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Symposium on Telescope Science

Editors:
Brian D. Warner
Dale Mais
David A. Kenyon
Jerry Foote

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Preface

The title says this is the 24th annual proceedings. The Society for Astronomical Sciences can trace its roots to the International Amateur-Professional Photoelectric Photometry (IAPPP), which was formed in June, 1980. The role of the IAPPP was to facilitate collaborative astronomical research between amateur, student, and professional astronomers, by providing a medium for the exchange of practical information not normally discussed at symposia or published in other journals.

The Western IAPPP Symposium, was held annually in the Southern California area starting in 1982. The symposium is held in Big Bear, California in the days before the RTMC Astronomy Expo. In 1998, the Western Wing of the IAPPP, was formed. In 2002, the Western Wing incorporated and in 2003 renamed itself the Society for Astronomical Sciences (SAS). The Society for Astronomical Sciences is a non-profit corporation exempt under I.R.S. Code Section 501(c)(3).

So, while under a new name, one thing that has not changed is the annual meeting, now called the Symposium on Telescope Science. Through this two-day event, the Society hopes to foster new friendships and new collaborations among amateur and professional astronomers with the goals being the personal scientific advancement of Society members, the development of the amateur-professional community, and promoting research that increases our understanding of the Universe.

It takes many people to have a successful conference, starting with the Conference Committee. This year the committee members are:

| Lee Snyder | Robert Stephens |
| Robert Gill | Dave Kenyon |
| Dale Mais | Brian D. Warner |
| Jerry Foote |

There are many others involved in a successful conference. The editors take time to note the many volunteers who put in considerable time and resources. We also thank the staff and management of the Northwoods Resort in Big Bear Lake, CA, for their efforts at accommodating the Society and our activities.

Membership dues alone do not cover the costs of the Society and annual conference. We owe a great debt of gratitude to our corporate sponsors: Sky and Telescope, Software Bisque, Santa Barbara Instruments Group, and Apogee Instruments, Inc.

Finally, there would be no conference without our speakers and poster presenters. We thank them for making the time to prepare and present the results of their research.

Brian D. Warner
Dale Mais
Dave Kenyon
Jerry Foote
Conference Sponsors

The conference organizers thank the following companies for their significant contributions and financial support. Without them, this conference would not be possible.

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The All Sky Camera Fireball Detector

David A. Kenyon
Professor of Astronomy, Sierra College
5000 Rocklin Rd.
Rocklin, CA. 95677
dkenyon@sierracollege.edu

Wayne T. Watson
Watson Laboratory
And James Clerk Maxwell Observatory
11467 Lothlorien Lane
Nevada City, CA 95959
Sierra_mtnview@earthlink.net

Abstract

A fireball (or bolide) is another name for a very bright meteor due to atmospheric entry of space debris, both natural and man-made. This paper will discuss the technology, methods and results from an “All-Sky” video camera and custom electronics to record fireball events that occur over Northern California. Images and links to videos will also be presented. The data was captured with Sandia Fireball Network cameras. These systems are operated by the Sierra College Astronomy Dept. and a private observatory in Nevada City, CA. These cameras are automated system which operates from just before dusk to just after dawn. The systems are video cameras, which have a 180° field of view. There is a custom signal processor and memory system, called the Sentinel system, which buffers a few seconds of continuous sky video. The processor compares each video frame to the previous, and when it detects an event (fireball), a few seconds of video (before and after the event) are sent to a host UNIX system. Only data relevant to motion is downloaded to the computer. The system provides a method of recording and study for meteor shower activity. Second, all events captured by both cameras provide the potential to compute the pre-earth-encounter orbit, but also to estimate the impact corridor of any meteorites the fireball might have produced. Our experience thus far shows that they occur about once a day and a very bright one every few weeks. The cameras began regular operation around mid-April 2004. To date, no fragments have been recovered from detected events. ©2005 Society for Astronomical Science.

1. Introduction

A fireball (or bolide) is another name for a very bright meteor due to atmospheric entry of space debris, both natural (meteoroid) and man-made.

Sandia National Labs in Albuquerque support a loosely connected network of meteor observers in the U.S. and Canada using video equipment provided to the observers. Sub-networks of individuals within the Sandia network work to coordinate visual, radio, and camera meteor data collected within the observer’s general geographic area. The only equipment not provided is a PC and Linux operating system. Our program represents two of the 6 (British Columbia, Nevada City, Rocklin, Albuquerque, Las Cruces, and El Paseo) observatories equipped with the latest cameras from Sandia. Older sites have video cameras and tape recorders, but no Sentinel software or electronics. Some sites have equipment developed by the owner.

There are 21 Canadian sites and 22 U.S. sites. The U.S. sites are west of the Rockies, mostly in Colorado, with others in Texas, New Mexico, Washington, and California. Canadian sites are spread from Newfoundland to British Columbia, i.e., coast to coast.

Figure 1. The Fireball Network
The most active programs are coordinated by Chris Peterson, who manages several sites in Colorado; Ed Majden, with his camera and a large network of observers in British Columbia; and Jim Gamble, a collection of observers in El Paseo.

Our experience thus far shows that fireballs occur about once a day and a very bright one every few weeks. Our program began regular operation around mid-April 2004. It is an automated system which operates from just before dusk to just after dawn.

2. Program Goals

Primarily, the system provides a method of recording and study for meteor shower activity. This program supports the Fireball Network as well as providing a focal point for meteor observations in the north western United States. It also provides a local educational resource as well.

Second, all events captured by both cameras provide the potential to compute the pre-earth-encounter orbit, as well as to estimate the impact corridor of any meteorites the fireball might have produced. This could result in the recovery of fragments from the falling object.

Since the beginning of the program over a year ago, we have received observer support and interest throughout California: with SAIC (for trajectory determination), UCLA (Astronomy Dept.), SETI/NASA (meteor studies), individual visual observers as well as several people who are planning to construct equipment.

3. The Hardware

The video system used is a Sandia National Laboratory Sentinel Video System. The system consists of a video camera with a wide field of view. The camera and lens are mounted pointed to the sky in a weather protected acrylic dome. This unit is secured to a stable platform with a clear view of the sky (see figure 1.). The system yields an “All-Sky” imaging system which has a 180° field of view.

The camera is connected (weather protected) to the Sentinel Video System (custom signal processor and memory) which buffers six seconds of continuous video of the sky. The processor compares each video frame to the previous frame and when it detects an event (fireball), six seconds of video (centered on the event) is sent to a host PC based Linux system.

The signal processor is connected to the PC via the printer parallel port. This port provides command and control interface as well as data transfer functionality.

There are two systems at different sites involved with this program. The first camera is located on top of the Sewell Hall Science Building on the Sierra College campus. Located at 38 deg. 47’ 25.9” of latitude and 121 deg. 12’ 49.3” longitude with an elevation of 90 meters in Rocklin, CA (USA), about 20 miles northeast of Sacramento, CA.
4. The Software

The Sentinel detection software is written in C and operates under the RT (realtime) Linux OS, which permits Sentinel to field events immediately and adjust data collection and related. Realtime OS’s are widely used in science and industry to optimize data collection under quickly varying conditions.

Clocks at both sites are independent, and provide time stamps for the data. The Rocklin site provides better accuracy and is used for analysis when both sites observe the same event.

The software provided with the Sentinel Video System includes control software (sentuser2) which is used to control the hardware. The software includes commands to initialize the system, set detection thresholds, masks to block problem areas of the sky and program on and off times.

Also provided with the system is a viewer (sdisplay) to examine the images and video clips downloaded from the hardware. There are also support programs to generate xxx.jpg [jpg image format] images and xxx.mov [Quicktime video formatted files] of detected events.

5. Data Collection

When an event is detected, an all-sky image is captured at 640 x 480 pixels of resolution and downloaded to the host. Superimposed on this background image is a 125 x 125 pixel video clip centered on the event (see figure 5). This image and video are viewable with the sdisplay program.
To accurately identify the path of the fireball event on the celestial sphere, an autoexpose feature is available on the *Sentinel System*. This feature periodically collects a long exposure image of the sky which yields the brighter stars for orientation of image and video clip (see figure 6).

The most useful data occurs when both of the *Sentinel Systems* in the program capture the same event. This provides the raw data to calculate the pre-atmospheric entry orbit and estimate the impact location. (see figure 7).

The coverage area for the program’s cameras is depicted in figure 8.

6. Preliminary Results

A summary of the 2004 events detected can be found in table 1.

![Figure 6. Autoexposure image.](image)

![Figure 7. A simultaneously detected event (Top: Rocklin, Bottom Nevada City).](image)

![Figure 8. Program coverage map.](image)

![Table 1: Counts for Both Sites](image)

An Example

On June 26th 2004, both sites recorded a very bright fireball (see figure 8). The images attracted quite a bit of attention in the meteor community. Initially the event was thought to be a Russian rocket reentry, but the event time and predicted re-entry were in disagreement. Robert Matson of SAIC decided to do an analysis of the images to extract sky positional elements.
From the two images and a calibration of the sky at each site, he was able to produce a 3-D track. It indicated a potential impact site southeast of Sacramento near Mosquito Ridge Road, a sparsely populated area in the Sierra Foothills sometimes used by sport motorcyclists. Robert decided there was enough potential to visit the area from July 17-18 to determine if a search was warranted. The terrain and estimated size of the fall indicated it did not merit putting together a search team. No fragments were found.

The program is being recognized as a primary contact for fireball observers in western U.S. Observers from Washington to California have contacted us for support, and others have provided visual information on sightings captured by the program cameras. We expect to continue to develop a network of observers on this coast for fireball sightings.

6. Detected Events

The images below are some of the brighter fireballs detected by this program.
8. Conclusion

The Program has made very good progress in establishing the two sites for fireball and meteor monitoring. We have been able to exchange information between the two sites and demonstrate that we can capture simultaneous events to predict a potential impact corridor. Working with other observers in the local area and throughout the west, we are becoming a focal point for meteor sightings in our local area.

For the future, we will continue to expand the program capabilities by improving upon the software and tool set which support the systems. Specific areas of focus will be: image stacking, operational ease, improve star overlays on images, and early estimates of tracks from dual observations of an event are some of the possibilities.

9. Acknowledgements

We would like to acknowledge the support provided by Richard Spalding and Joseph Chevez with the Sentinel hardware and software. Also, the support of Sierra College student Chris Giorgi for his Linux support of the systems.

10. References


DK Canum Venaticorum: A Dynamic Eclipsing Binary Star

Robert A. Koff
Antelope Hills Observatory
980 Antelope Drive West
Bennett, CO 80201
bob.koff@worldnet.att.net

Dirk Terrell
Department of Space Studies, Southwest Research Institute
1050 Walnut Street, Suite 400
Boulder, CO 80302
terrell@boulder.swri.edu

Arne A. Henden
Universities Space Research Association/US Naval Observatory, Flagstaff Station
P. O. Box 1149
Flagstaff, AZ 86002-1149
aah@nofs.navy.mil

Timothy Hager
Mount Tom Observatory
34 Mount Tom Road
Milford, CT 06776
thager6164@earthlink.net

ABSTRACT:
A study undertaken to characterize the eclipsing binary system DK CVn revealed a very unusual lightcurve. During follow-up observations over the next three years, the system was observed to be undergoing significant changes. In 2005, more observations are being made, and preliminary interpretation of the system is being undertaken. © 2005 Society for Astronomical Science.

1. The Observational Campaign

DK CVn was originally identified as a suspected eclipsing binary star by the ROTSE survey (Akerlof et al, 2000). Times of minimum and basic parameters of the system were published in 2001, along with the remark, “pronounced reflection effect” (Die-thelm, 2001). The AAVSO EB Team selected it as a potential target for study in 2002. Observations by Koff revealed that the system was exhibiting unusual behavior. See figure 1.

Only the data from Henden and Hager are shown, for clarity. The phase diagram is based on a period of 0.49496 day. Terrell noted that it resembled the V361 Lyr system, which was itself unique. Notice the large increase in amplitude following the primary eclipse, as compared to the magnitude preceding the primary eclipse. This behavior is what distinguishes this system from all others except for V361 Lyr.

Observing intensified during the remainder of the 2002 observing season. Henden carried out time series photometry of the star in UBVRI bands, and a number of amateur astronomers using BVRI filters also contributed to the campaign.
The following year, 2003, follow-up observations by Koff and Howell indicated that the star system had undergone changes. See figure 2.

Here, the shape of the lightcurve following the primary eclipse has changed. There is also evidence that the primary and secondary eclipses have deepened.

This led to further observations in 2004 by Koff, which revealed yet more changes in the lightcurve. See Figure 3.

There was significant change between 2003 and 2004. Note how the magnitude hump following the primary eclipse has almost leveled off with the ingress shoulder value. The primary eclipse has deepened, while the secondary eclipse has become shallower. At this apparition, the curve resembles that of a more typical eclipsing binary system.

Notice also that the magnitude values at the 75% phase position have remained unchanged throughout the period of the study. At this point in the lightcurve, the two stars are both visible, side by side.

In 2005, Koff began observing the system in January, early in the observing season. It was immediately apparent that the lightcurve was again showing unusual behavior, similar to that of 2002. See figure 4.
2. Acknowledgments

The authors wish to express appreciation to the AAVSO Eclipsing Binary Team observers who assisted in this project, but whose observations are not included in this report.

3. References


Diethelm, R. Times of Minimum for Eclipsing Binaries from ROTSE1 CCD Data, III: Variables Classified as Type E. IBVS 5060
HD23642 and the Distance to the Pleiades

Dirk Terrell
Southwest Research Institute
1050 Walnut St., Suite 400
Boulder, CO 80302
terrell@boulder.swri.edu

Abstract
I discuss the nature of the newly discovered eclipsing binary HD 23642 and its role in settling the disagreement between the Hipparcos distance to the Pleiades cluster and the distance found by other methods. © 2005 Society for Astronomical Science.

1. Introduction
Open and globular clusters have played a crucial role in the development of stellar evolution theory. When we observe a cluster, we are observing a group of stars that have approximately the same distance (since the size of the cluster is usually small compared to its distance), age (since star formation happens fairly quickly once it starts) and chemical composition (since the stars all formed out of the same cloud of dust and gas). These properties enable us to test various predictions of stellar evolution theory that are virtually impossible to do for single stars. For example, theory predicts that more massive stars should evolve quickly while low mass stars should evolve very slowly. Observations of clusters show that this is exactly what happens. You can plot brightness versus color for the stars in a cluster and you will find that the hotter (bluer) stars are moving away from the main sequence while the cooler (redder) ones remain on the main sequence.

Clusters thus play an important role in our understanding of the structure and evolution. Therefore it was quite disconcerting when the parallax observations of Hipparcos seemed to indicate that the distance to the Pleiades was about 10% smaller than previous determinations, implying that theoretical stellar models of the stars were about 0.3 magnitudes too bright. There was a scramble to explain how this discrepancy might arise in terms of such things as an unusual chemical composition of Pleiades stars, but none of these explanations withstood the test of observational scrutiny. Perhaps the Hipparcos distance was in error?

It is now clear that the Hipparcos distance was indeed in error due to a subtle problem with the reduction of the Hipparcos observations. Recent measurements of two binary star systems, the spectroscopic/interferometric binary Atlas (HD 23850) and the spectroscopic/eclipsing binary HD 23642, show that the distance to the Pleiades is approximately 132 pc rather than 116 pc as found by Hipparcos (van Leeuwen, F & Hansen Ruiz 1997. Zwahlen, et al., (2004) presented radial velocities and interferometric observations of Atlas and found the distance to be 132±4 pc. Munari, et al. (2004) analyzed radial velocities and BV photometry of HD 23642 and found a distance of 132±2 pc. Parallax measurements using the Fine Guidance Sensors on the Hubble Space Telescope (Soderblom, et al. 2005) result in a distance of 134.6±3.1 pc. Percival, et al. (2004) use optical photometry and infrared data from the Two Micron All Sky Survey to estimate the Pleiades distance as 133.8 ±3 pc.

2. HD 23642
In collaboration with G. Torres, discover of the eclipsing nature of HD 23642 (Torres, 2003), C.H.S. Lacy, L. Marschall, and I. Ribas, I have analyzed UBVI photometry and radial velocities of HD 23642 with the goal of accurately determining the distance to the system. As seen in Figures 1 and 2, the eclipses are very shallow and require many high precision observations for accurate measurements. The system is detached and the eclipses are partial, presenting some challenges for the analysis, such as the severe correlation between the ratio of the stellar radii and the inclination, that must be dealt with if the results of the solution are to be believed. At press time, the analysis is still in progress but I will discuss our analysis and results at the Symposium.
3. Acknowledgements

This project was supported in part by a grant from the American Astronomical Society’s Small Research Grants Program.

4. References


Torres, G. 2003, Discovery of a Bright Eclipsing Binary in the Pleiades Cluster, IBVS #5402.


**Fundamentals of Solving Eclipsing Binary Light Curves Using Binary Maker 3**

David H. Bradstreet  
Eastern University  
Department of Physical Science  
1300 Eagle Road  
St. Davids, PA 19087-3696  
dbradstr@eastern.edu

**Abstract**  
The parameters and fundamentals of eclipsing binary light curve analysis are discussed and strategies presented for solving them using Binary Maker 3. The major types of eclipsing binaries are discussed in some detail so as to give users a starting point in their analysis. The online Catalog and Atlas of Eclipsing Binaries (CALEB – http://caleb.eastern.edu/) is also briefly discussed. © 2005 Society for Astronomical Science.

**1. Introduction**  
The study of binaries is the only way to directly measure the masses of stars, arguably the most important characteristic of stars that we need in order to understand them. Eclipsing binaries, binaries whose orbital plane nearly lies in the observer’s line of sight, offer us fantastic astrophysical laboratories through which we measure not only their masses, but their temperatures, sizes, luminosities, and other intrinsic properties of the stars. Understanding and analyzing the way the light of the two stars changes as they rotate and revolve, when coupled with spectroscopic data revealing their periodic radial velocities, can unlock the secrets of these systems. The knowledge mined becomes the bedrock upon which stellar astrophysical theories are tested.

The first part of this paper is devoted to defining and explaining the many input parameters needed to construct meaningful eclipsing binary models. The second part contains strategies for recognizing various types of eclipsing binary and implementing those strategies to successfully model light curves. Hopefully misconceptions and mysteries related to the field of binary star research and analysis will be clarified, and more people will be emboldened to enter this fascinating aspect of astrophysics. Because of the limitations of space, we will limit this discussion to circular orbit systems.

There are several computer codes available which can analyze eclipsing binary light curves, but the one which will be used throughout this paper is Binary Maker 3 (Bradstreet and Steelman (2002)), a user-friendly Java program that runs on Windows, Macintosh, Linux and Sun Solaris (Unix) platforms. All of the graphics in this paper as well as the analysis were created using this program. More information concerning the program can be found at www.binarymaker.com. The main windows of Binary Maker 3 are shown in Figure 1. The upper left window allows the user to enter the necessary parameters for constructing the binary, including star-spot parameters. The upper right window displays a 3D model of the system as it revolves and rotates. The lower left window plots the observed and theoretical light curves of the binary. The lower right window shows the observed and theoretical radial velocity curves of the binary.

**2. Binary Maker 3 Parameters Explained**  
The following sections are devoted to defining and explaining the major input parameters needed to construct meaningful eclipsing binary models. Although specific to Binary Maker 3, most of them are common to other light curve synthesis programs.

**2.1 Grid Values**  
In order to mathematically represent the stars and the light which they emit, surface grid points are created using Roche equipotentials. These surfaces, where the gravitational potentials are the same (hence the name equipotential) take into account the masses of each star, their separations, and their rotation and revolution. At each grid point the temperature, black body emission, limb darkening, visibility, etc., are calculated throughout the system’s orbit in order to recreate its light curve.
The *latitude grid number* is the number of surface area elements to be calculated from the pole of the star to its equator. The *longitude grid number* is the number of surface area elements to be calculated along the major (largest) axis of the star along its equator. To make the number of *longitude elements* comparable to the *latitude elements*, you should make this number twice as large as the *latitude grid number*, in this case 40 if the latitude grid value is 20. A few examples are shown below in Figures 2-4.

Figure 1. Four main windows of Binary Maker 3 showing AH Aur

Figure 2. AB And displayed at phase 0.25P at 10 x 20 grid size

Figure 3. AB And displayed at phase 0.25P at 20 x 40 grid size

Figure 4. AB And displayed at phase 0.25P at 40 x 80 grid size
The larger the grid values the more accurate will be the modeling, but the computation times will rise geometrically.

2.2 Mass Ratio (q)

The mass ratio \( q \) is usually defined as the mass of the less massive star (\( M_2 \)) divided by that of the more massive star (\( M_1 \)), i.e.,

\[
q = \frac{M_2}{M_1}
\]

Therefore this quantity is usually less than 1.00. However, if the less massive star is hotter than the more massive component, then the deeper eclipse (by convention the primary eclipse) will occur when the less massive star is eclipsed. In order to shift the phases by 0.50 to account for this, the mass ratio is inverted, as for W-type W UMa systems. In Binary Maker 3 this convention is observed, i.e., for hotter secondaries you should input the inverse mass ratio, i.e., greater than 1.0, but star 1 is still considered to be the more massive star and star 2 the less massive one.

2.3 Modified Omega Potential (\( \Omega \))

The modified equipotential \( \Omega \), along with a specified mass ratio, completely describes the surface structure for synchronously rotating, circular orbit binary stars. It takes into account both the gravitational and centrifugal forces in the binary. As the name implies, a gravitational equipotential demarcates the surface along which the gravitational potential energy is constant. Thus the stars will take the shape of such an equipotential, just like the Sun is almost a sphere because that shape is a gravitational equipotential for a single star that is rotating very slowly. As the value of the equipotential is increased, the size of the star decreases. This makes sense since the gravitational potential increases as one approaches the mass center. \( \Omega_{\text{inner}} \) is the value of the inner critical Roche equipotential and represents the point at which the stars just come into contact. \( \Omega_{\text{outer}} \) is the value of the outer critical Roche equipotential and represents the limit to stability to any overcontact system since the outer potential has a "hole" in it (gravitational acceleration = 0) and gas will leave the system. Representative equipotentials for a mass ratio \( q = 0.40 \) are shown in cross-section in Figure 5.

\[
\Omega_{\text{inner}} \leq \Omega < \Omega_{\text{outer}}
\]

The fillout factor \( f \) for detached stars will lie between \(-1 < f \leq 0 \). The fillout factor \( f \) for overcontact systems will lie between \( 0 \leq f \leq 1 \) i.e., from in contact with the inner critical surface (\( f = 0 \)) to being in contact with the outer critical surface (\( f = 1 \)). Several examples of binary systems with different fillout factors (all with mass ratios equal to 0.40) are shown below in Figures 6-10, along with their inner and outer critical Lagrangian surfaces.

Another definition of fillout factor or parameter was set forth by Mochnacki and Doughty (1972) and is shown below:

\[
F = \frac{\Omega_{\text{inner}} - \Omega}{\Omega_{\text{inner}} - \Omega_{\text{outer}}} + 1 \quad \text{for } \Omega < \Omega_{\text{inner}}
\]

\[
F = \frac{\Omega_{\text{inner}}}{\Omega} \quad \text{for } \Omega > \Omega_{\text{inner}}
\]
In this case the fillout $F$ for overcontact stars lies between $1 \leq F \leq 2$.

Rucinski (1973) defined yet another fillout factor similar to the one used in *Binary Maker 3*:

$$f_{\text{Rucinski}} = \frac{\Omega - \Omega_{\text{outer}}}{\Omega_{\text{inner}} - \Omega_{\text{outer}}}$$

In this definition of fillout the values would be $f_{\text{Rucinski}} = 0.0$ for the stars in contact with their outer Lagrangian surfaces, and $f_{\text{Rucinski}} = 1.0$ for the stars just filling their inner contact surface, *i.e.*, just in contact with each other at the inner Lagrangian point $L_1$. The relationships between the three fillout parameters are as follows:

$$f = F - 1$$

$$f_{\text{Rucinski}} = 2 - F = 1 - f$$

### 2.5 Fractional Radii

The distance between the mass centers of the two stars is defined to be unity for circular orbits. *Binary Maker 3* allows for specifying equipotential surfaces by the input of the fractional stellar radius $a$ or $r_{\text{back}}$ in Wilson-Devinney notation (Wilson and Devinney (1971)). This radius is along the line of centers between the two stars and is the radius that is always facing away from its companion. This parameter cannot be used when specifying contact or overcontact systems since the input radii would have to be perfect in order to generate exactly the same potential for both stars. All of the commonly specified stellar radii are depicted in Figure 11.

### 2.6 Effective Wavelength

The wavelength parameter refers to the effective wavelength in angstroms (Å) of the filter used to acquire the data. The default value is 5500 Å corresponding to the standard Johnson wideband $V$ filter. Other effective wavelengths for typical photometric systems are given in Tables 1-2.

<table>
<thead>
<tr>
<th>Johnson</th>
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<tbody>
<tr>
<td>$U = 3600$ Å</td>
</tr>
<tr>
<td>$B = 4400$ Å</td>
</tr>
<tr>
<td>$V = 5500$ Å</td>
</tr>
<tr>
<td>$R = 7000$ Å</td>
</tr>
<tr>
<td>$I = 8800$ Å</td>
</tr>
<tr>
<td>$R_{\text{Kron-Cousins}} = 6400$ Å</td>
</tr>
<tr>
<td>$I_{\text{Kron-Cousins}} = 7900$ Å</td>
</tr>
</tbody>
</table>

Table 1. Johnson effective wavelengths
Figure 11. Definitions of radii; diagram seen from the plane of the binary orbit

2.7 Temperature

The temperature parameter refers to the mean surface effective temperature of the stars as discussed by Wilson (1979).

2.8 Gravity Brightening (Darkening) Exponent

Von Zeipel (1924) proved that for totally radiative stars the surface flux was directly proportional to the value of the gravitational acceleration \( g \) at the stellar surface. (The gravitational acceleration is \(- \nabla \Omega \)). The equation used to determine the local temperature \( T_{\text{local}} \) on the surface of the stars is:

\[
T_{\text{local}} = T_{\text{eff}} \left( \frac{g}{g_n} \right)^{0.25}
\]

where \( T_{\text{eff}} \) is the mean surface effective temperature of the star, \( g \) is the local gravity at a specific surface area element and \( g_n \) is the value of \( g \) at the poles. The Gravity Brightening Exponent \( \alpha \) is 1.00 for radiative stars and according to Lucy (1967) is 0.32 for convective stars. It is also wavelength independent. The approximate temperature demarcation for radiative and convective stars is 7200 °K, i.e., stars hotter than 7200 °K are believed to have radiative envelopes, and stars cooler than 7200 °K will possess convective envelopes.

2.9 Limb Darkening (wavelength dependent)

Limb darkening is the diminishing of the brightness of a star’s surface as one looks away from its center, towards the edge of the disk (or limb). This is caused by the fact that stars are hotter the deeper into them you go, and when you are looking towards the center of their disk, your line of sight protrudes most deeply into their photospheres. When you look towards the edge of the disk your line of sight does not venture as deeply below the surface of the star, and hence the temperature you see is less and the gas gives off less light.

**Binary Maker 3** uses the linear limb darkening law. The coefficient \( x \) is used in the linear limb darkening equation

\[
I(\theta) = I(0^\circ)(1 - x + x \cos \theta)
\]
where $\theta$ is the angle measured between your line of sight and the surface normal of a particular surface area element. When you are looking straight at the center of the star $\theta = 0^\circ$ and the outer limb of the disk is at $\theta = 90^\circ$. These values are wavelength and temperature dependent and theoretical values must be looked up in tables because most light curves do not yield trustworthy values. The table of Al Naimiy (1978) contains the values most commonly cited prior to 1993. The most commonly used values since 1993 are those of Van Hamme (1993). Both tables of limb darkening coefficients are also included within the Help files of Binary Maker 3 as well as in the appendices of the User Manual. The tables give various limb darkening coefficients for different wavelengths and the log of the gravitational acceleration at the surface of the star. A $\log (g) \approx 4$ is typical for most stars.

### 2.10 Bolometric Albedo (Reflection Coefficient)

When the radiation from one star strikes the surface of the other, its energy will heat up the receiving surface and, if the star is mostly radiative, re-radiate the energy like a mirror. The bolometric albedo (sometimes called the reflection coefficient) is the percentage of incident radiation that is re-radiated by the companion star. For radiative stars the coefficient is assumed to be 1.00. Rucinski (1969) showed that for convective stars the albedo will be approximately 0.50 because surface convection will carry away some of the energy to re-radiate it from regions other than where it was incident. Thus for stars whose temperatures are less than 7200 °K this coefficient should be set to 0.50.

### 2.11 Luminosity (wavelength dependent)

The luminosity of a star is a measure of its total energy output. Binary Maker 3 outputs the luminosities $L_1$ and $L_2$ that indicate the percentage of the total luminosity each star emits. The total luminosity is normalized to 1.00, so that $L_1 + L_2 = 1.00$. These quantities are wavelength dependent.

### 2.12 Third Light (wavelength dependent)

This wavelength dependent parameter takes into account any contamination in the data due to sources other than the binary stars themselves. This can be a third star that is part of the binary or even a field star that could not be excluded from the diaphragm when observing at the telescope because of its coincidental proximity to the binary. Third light is in the same units as the input or output flux and is not a luminosity fraction like $L_1$ and $L_2$.

### 2.13 Inclination

The inclination of a binary is a description in degrees of the tilt of its orbital plane with respect to an Earth observer. Therefore, as shown in Figure 12, an inclination of $0^\circ$ corresponds to looking down onto the poles of the stars. An inclination of $90^\circ$ means that the orbital plane of the two stars lies directly in the Earth observer's line of sight; thus eclipses of some kind are guaranteed. Figure 12 shows the semi-detached system AI Cru at several inclinations of varying degrees.

![Figure 12. AI Cru viewed at different inclination angles](image)

### 2.14 Orbital Phase and Normalization Phase

The orbital phase of a binary is simply a means of describing the positions of the stars at any point within their orbit. The phase varies from 0.00 to 1.00, and is a fraction of the period ($P$) of the binary. The deeper eclipse (primary eclipse = hotter star eclipsed) is defined to be at phase 0.00P, and secondary eclipse (secondary eclipse = cooler star eclipsed) for circular orbits is half way through the cycle or 0.50P. The quadrature phases (when both stars are most visible) are naturally in between these two eclipse phases for circular orbits and are specified as 0.25P and 0.75P phase. These phase orientations are shown below in Figure 13.

![Figure 13. MR Cyg shown at four phase orientations](image)

The Normalization Phase is the phase at which you wish to exactly pin your synthetic light curve to your observed one. Any phase can be chosen, but usually one of the quadratures (0.25P or 0.75P) is chosen. This is entirely up to the user.

### 2.15 Normalization Factor

The Normalization Factor is the actual value that you wish for your synthetic light curve to have at the chosen Normalization Phase. Thus this parameter allows you to match your theoretical light curve exactly to your observed one (at least at that one phase). Usually when a light curve is prepared for analysis its magnitude values (logarithmic scale) are...
converted into flux values (linear scale) and the highest value of the flux (usually at one of the quadrature phases 0.25P or 0.75P) is defined to be unity (1.00). As with the Normalization Phase, the value of the Normalization Factor is entirely up to the user and depends on how the user set up their observed light curve for analysis.

3. The Major Close Binary Types

So, how does one go about solving or analyzing an eclipsing binary light curve? The first step is to become familiar with the fundamental types of eclipsing binaries that exist, and recognizing the general morphology that each of their light curves exhibits. The overall classification of eclipsing binary systems is founded upon Roche equipotentials. The shapes of the two stars will be mandated by their mass ratios, rotation rates and their separations. The different shapes that are possible for a mass ratio of 0.40 (synchronous rotation, circular orbit) were shown above in the section describing the modified Omega potential (see Figure 5), where the configuration (shapes of the stars’ surfaces) depends upon how close the stars reside to their inner Lagrangian surface.

There are basically six kinds of eclipsing binary systems:
1. detached
2. semi-detached
3. near contact
4. contact
5. overcontact
6. double contact

These are discussed in some detail below.

3.1 Detached Systems

Detached systems occur when both stars are located well within their respective inner Lagrangian surface and therefore take on essentially spherical shapes, as shown below for KP Aql in Figure 14.

![Figure 14. Inner and Outer Lagrangian surfaces for the detached binary KP Aql](image)

3.2 Semi-Detached Systems

In larger mass binaries, the star with the larger mass will evolve more rapidly than its lower mass companion, and as it expands and fills its inner Lagrangian surface it will spill mass onto its companion through its inner Lagrangian point L1. As the mass transfers over to the less massive star, the mass ratio changes (increases) in such a way that the inner Lagrangian surface for the more massive component decreases in size, accelerating the mass transfer over to the less massive star. Eventually the more massive component becomes the less massive component and we are left with a star that has filled its inner Lagrangian surface through stellar evolution but is now the less massive component. The now more massive component is the less evolved star, and this apparent mystery has been dubbed the Algol Paradox. AD Her is a typical Algol system (or semi-detached system) and is shown in Figure 15.

![Figure 15. Inner and Outer Lagrangian surfaces for the semi-detached binary AD Her](image)

3.3 Near Contact Systems

Near contact binaries are a bit more loosely defined. Strictly speaking these are just more extreme examples of semi-detached systems where one component is in contact with its inner Lagrangian surface and the other is nearly in contact. The accepted paradigm for these systems is that they are approaching the contact stage, where both stars fill their inner Lagrangian surfaces. V1010 Oph, a typical near contact system, is shown in Figure 16.

![Figure 16. Inner and Outer Lagrangian surfaces for the near contact binary V1010 Oph](image)

3.4 Contact Systems

Contact binaries are those systems wherein both stars exactly fill their inner Lagrangian surfaces. It is believed that these systems are in the stage between near contact and overcontact, and some theories be-
believe that they will oscillate between these two configurations until they finally become overcontact systems, possibly on their way to finally coalescing into a single star. A typical contact system, BX And, is shown in Figure 17.

![Figure 17. Inner and Outer Lagrangian surfaces for the contact binary BX And](image)

### 3.5 Overcontact Systems

Overcontact binaries are those systems wherein both stars overflow their inner Lagrangian surfaces. The stars are nearly at the same temperature, usually within a few hundred degrees Kelvin, the smaller star stealing energy from its more massive companion. A typical overcontact binary, AE Phe, is shown in Figure 18.

![Figure 18. Inner and Outer Lagrangian surfaces for the overcontact binary AE Phe](image)

### 3.6 Double Contact Systems

Double contact binaries consist of stars where both systems exactly fill their inner Lagrangian surfaces but one or both of the components are rotating faster than synchronously with their revolutionary period. These systems appear to be relatively rare, but they do exist. The best known example of these interesting binaries, RZ Sct, is shown in Figure 19.

![Figure 19. Inner Lagrangian surfaces for the double contact binary RZ Sct](image)

### 4. Recognizing Light Curves of Different Binary Types

If you become familiar with the different shapes of eclipsing binary light curves, you will usually be able to quickly hone in on a reasonable solution. The key is in understanding why different types of binaries exhibit different shaped light curves, as well as understanding exactly what each light curve parameter does in the creating of a synthetic light curve. We will begin with showing sample light curves of the various binary types listed above and explaining why the light curves appear as they do. We will also limit our discussion in this paper to circular orbit systems.

#### 4.1 Detached Eclipsing Binaries

Typical detached binaries consist of relatively spherical stars, and hence as they revolve about each other the light level remains fairly constant and so the out of eclipse light variation will be quite small. Let us look at the b (Strömgren blue) light curve of GZ CMa from Andersen et al. (1985), shown below in Figure 20.

![Figure 20. b (Strömgren blue) light curve of GZ CMa from Andersen et al. (1985)](image)

The out of eclipse light level is indeed mostly constant, and the eclipses are well-defined and sharp, indicating that the stars eclipse each other partially or that the stars are the same size and shape (spheres). The difference in depths is mostly due to the fact that the stars must possess different surface temperatures. This makes sense because, for circular orbit systems, the amount of surface area eclipsed at both eclipses must be the same. Therefore, since the same area of star’s surface is blocked at the eclipses, the difference in light loss must be primarily due to a difference in surface temperature. In other words, when the hotter component is eclipsed, you will lose more light than when the cooler component is eclipsed. So, in the case of GZ CMa, the deeper eclipse is caused when the cooler star blocks the hotter star. The width of the eclipses is a function of the sizes of the stars as well
as their orbital inclination. The greater the inclination the wider and deeper the eclipses will be.

To begin the analysis of a light curve, one needs as much information about the binary as possible. Especially useful is a color index or spectroscopic classification because either will allow an approximate temperature designation for the stars. Once an approximate temperature has been established, limb darkening coefficients can also be determined from the appropriate tables for the effective wavelength of the observations. Prior to 1993 the tables published by Al-Naimy (1979) were widely used and they are included in the Appendix of the Binary Maker 3 User Manual and in the Help files. Van Hamme (1993) published completely revised modern limb darkening coefficients based upon Kurucz (1989) model atmospheres. A portion of these tables is included in Appendix of the User Manual and Help files of Binary Maker 3 through the kind permission of Dr. Van Hamme.

Perhaps the most elusive parameter is the mass ratio. For partially eclipsing systems the direct determination of this parameter is quite problematic because many different mass ratios can be used with various other parameters to mimic the observed light curves. A spectroscopic mass ratio is needed in order to pin down the mass ratio of partially eclipsing systems, as well as to allow the direct calculation of the absolute parameters of the binary (masses, radii, etc.). If you do not have a spectroscopically determined mass ratio but the system exhibits a total eclipse, then it is possible to determine a fairly reliable mass ratio because of the restrictions inherent in a total eclipse (i.e., the length of totality puts severe constraints on the relative sizes of the stars).

If the binary has no radial velocity curve and is partially eclipsing, then the best one can do is to make a series of models with a range of mass ratios (q) and plot the residuals for each fit versus the mass ratio used for that particular fit. This technique, sometimes called the “q-method,” is certainly not guaranteed to determine the true mass ratio, but it is the best that can be done in the absence of more information. Certainly for detached systems like GZ CMa the mass ratio cannot be too far from unity since the stars are spherical and about the same temperature with fairly deep eclipses. Fortunately GZ CMa has an excellent radial velocity curve (Popper et al. (1985)) and a well-determined spectroscopic mass ratio of 0.909. Armed with this and color indices indicating surface temperatures close to 8500 °K, one can quickly converge to an excellent model for this system by trying different radii and inclinations until the light curve is closely approximated. The solution given by Popper et al. (1985) is shown below in Figure 21.

The final parameters for GZ CMa can be found in the Zip folder of Binary Maker 3 as well as on the Catalog and Atlas of Eclipsing Binaries (CALEB) at http://caleb.eastern.edu/. The relative sizes of the stars and their inner and outer critical Lagrangian surfaces are shown below in Figure 22. Note that the stars are well within their inner Lagrangian surface and therefore they are nearly spherical, hence the lack of out of eclipse light variation.

Let us explore another detached binary, EE Peg, as shown in Figure 23 from Ebbighausen (1971). This system is not only obviously detached (as evidenced by its relatively flat out of eclipse light curve) but its components possess very different surface temperatures as seen from the large difference in eclipse depths. The secondary eclipse is also total, which again greatly constrains the model.

31
These stars are also mostly spherical, but not as much as GZ CMa, and the out of eclipse light variation is more evident than in the light curve of GZ CMa. The stars of EE Peg are shown in Figure 24 with respect to their Lagrangian critical surfaces. Note that the larger, more massive and hotter star is slightly more ellipsoidal than the larger component of GZ CMa and causes most of the out of eclipse light variation seen in the light curve.

If the stars are closer together and/or their surfaces are in close proximity to the inner Lagrangian surface, their shapes will begin to distort roughly into ellipsoids because of gravitational tidal forces. Thus, because the stars are “football” shaped, their light curves will vary continually even when not eclipsing because their visible cross-sectional areas will be continually changing. We expect the out of eclipse portions of the light curve to not be flat but varying in flux, as seen in the close binary system NN Cep (Gudur et al. (1983)) whose light curve is shown in Figure 25.

The model of NN Cep is shown in Figure 26 where the larger component is fairly close to its inner critical surface and the smaller component is also close and therefore both stars are distorted from spheres and create varying light levels even out of eclipse.

Another system where both stars are very close to their inner critical surfaces is AI Cru (Bell et al. (1987)), whose light curve is shown in Figure 27 and the model of the system itself is shown in Figure 28.
Note that the light curve of AI Cru is reminiscent of NN Cep except that the eclipse depths are more dissimilar in AI Cru and the widths of the eclipses are wider. Thus the stars of AI Cru must be both fairly different in temperature (different eclipse depths) as well as much larger stars relative to each other (to explain the longer eclipse durations).

4.2 Semi-Detached Eclipsing Binaries

The origin of semi-detached (Algol) systems was discussed already in section 3.2. A typical semi-detached system, AS Eri (Koch (1960); Van Hamme and Wilson (1984)), is shown in Figure 29 and its critical surfaces and model are given in Figure 30.

4.3 Near-Contact Eclipsing Binaries

When one star is in contact with its inner critical surface and its companion is nearly in contact with its critical surface, the binary has been described as a near contact system. The light curve changes in magnitude throughout the orbit but the stars are usually quite different in temperature, as can be seen in the near contact system AK CMi (Samec et al. (1998)) shown in Figures 31 and 32.

4.4 Contact and Overcontact Eclipsing Binaries

The next possibility is the contact system, when two stars both exactly fill their inner Lagrangian surfaces. The stars typically possess considerably different temperatures, and of course because they are very much distorted from spheres, the light levels vary throughout the entire orbit. It is often difficult to distinguish between a near contact system and a contact system, and careful light curve analysis must be used to discern which type the binary is. Usually the temperatures of the stars are not quite as disparate as in a near contact system, probably because the stars are in a stage where the smaller star is beginning to steal energy from the hotter, more massive component as they come into contact. A good example of a contact system is BX And (Samec et al. (1989)), shown in Figure 33 and its model is given in Figure 34.
A careful comparison of the light curves of AK CMi (near contact) and BX And will reveal more similarities than differences, the main difference being the greater depth of the secondary eclipse demonstrating that the secondary (less massive, smaller, cooler) star is more nearly the same temperature as the primary (more massive, larger, hotter) star.

It is believed that the stage after contact systems is the overcontact binary. These stars have been in contact for a long enough period of time such that their surfaces have come into an approximate thermal equilibrium and their surfaces have extended beyond their inner critical surfaces. This results in light curves that are continually changing in the flux level and the eclipse depths are nearly equal because of the nearly equal surface temperatures. Shown in Figure 35 is a typical overcontact binary (fillout = 0.14), AD Cnc (Samec 1989), which demonstrates the characteristics of this type of light curve.

If the shoulders of the eclipses are steeper than AD Cnc, then this is an indication that more light is being cut off more quickly, and this often means that the stars are even more extended beyond their inner critical surfaces, i.e., their fillout is larger than the average value of 0.15 for most overcontact systems (Robertson and Eggleton 1977). This is shown in Figure 36 for AW Lac (Jiang et al. (1983)), an overcontact system with a large fillout = 0.60.

Figure 37 shows synthetic light curves for a typical overcontact binary but for six different values of fillout. The top curve represents a fillout = 0.00 demonstrating broad shoulders and the slowest drop-off in light level. Each successive curve represents an increase in fillout of 0.20 (i.e., 0.20, 0.40, 0.60, 0.80 and 1.00) up to the maximum allowed of f = 1.00. It is easily seen that as the stars increase in size they block off more light and at an accelerated pace as evidenced by the steep eclipse shoulders as the fillout increases. No other binary parameters were changed other than the value of the fillout; the inclination was not changed.
Figure 37. The effect of varying the value of the fillout on the shape of an Overcontact binary light curve. The uppermost curve shows the system with a fillout of 0.00, and each successive curve beneath it has a larger fillout by increments of 0.20 (i.e., 0.20, 0.40, 0.60, 0.80 and 1.00).

5. Light Curve Analysis “Tips”

An excellent way to get a “feeling” for what kind of parameters will create what kinds of light curves is to explore the Catalog and Atlas of Eclipsing Binaries at http://caleb.eastern.edu/. You can list the hundreds of on-line binaries by binary type and look at the light curves and perhaps find one that is similar to your own. Many of the systems listed in this database are also included in the Zip library accompanying Binary Maker 3. Additional systems can be downloaded directly from the site.

Once you have determined the basic type of binary from the shape of the light curve, there are several straightforward techniques that can be used to hone in on a possible solution.

If the light curve appears to be an overcontact system (i.e., relatively short period (less than 1 day), light variation throughout the light curve and similar depth eclipses) then put the program into fillout mode. If color indices exist then an approximate effective temperature can be set for both stars, i.e., initially set both stars to the same temperature. If the period is less than 0.4 days, then the system will likely be a W-type overcontact system where the larger star is the cooler star. This requires a mass ratio input that is greater than unity. Obviously use a spectroscopic mass ratio if that is available. If the system displays a total eclipse then the mass ratio can be determined through trial and error modeling of different mass ratios until the eclipses are properly fitted. If the system is partially eclipsing than a series of solutions with different mass ratios will have to be attempted, and the best-fit parameters will be the best estimate of the actual ones.

Once a mass ratio has been input, a good starting point for both fillouts is 0.15, the average value found for overcontact systems (Robertson and Eggleton (1977)). Start with an orbital inclination of 85° unless the system exhibits total eclipses. If it displays total eclipses then begin with 90°. If the estimated temperatures are less than 7200° K, then treat the stars as convective and let the gravity brightening coefficients be 0.32. If the stars are believed to be higher in temperature than 7200° K (mostly radiative stars), then set the gravity brightening coefficients be 1.00. Linear limb darkening coefficients can be determined from Van Hamme’s (1993) tables that are available in the program Help files as well as in the Appendix of the User Manual. The reflection coefficients will be 0.5 for convective stars and 1.0 for radiative stars.

Assuming that the light curve has been read into the Light Curve Plot and that you have generated a theoretical (synthetic) model, compare the theoretical light curve to the observations. If the theoretical eclipses are both too deep, then decrease the inclination until the depths are more comparable to the observations. If one eclipse seems the correct depth but the other is not, then the temperature of one of the stars must be changed (usually the smaller star is changed). Light curve analysis does not give the absolute temperatures of stars but rather the temperature difference between the two stars. Therefore changing the temperature of just one of the stars is sufficient. If the eclipse that doesn’t fit is too low then the star being eclipsed is too high in temperature. If the theoretical eclipse isn’t low enough, then that star is too cool. Note that adjusting the temperature of one star will affect the relative depths of both eclipses. When both eclipses are too low or two high from the observations by approximately the same amount, then simply adjust the inclination at that point to make both the eclipses fit.

Once a solution for a particular wavelength has been achieved, save it as a BM3 file (input parameter file for Binary Maker 3) and then load in another bandpass dataset and change the wavelength and limb darkening coefficients to match that wavelength (also third light if it is nonzero). If your first model is a good one, then the other wavelength-independent parameters should fit the other dataset as well. If it does not, then adjustments will have to be made in the model (temperatures, sizes, mass ratio, etc.) until a coherent model satisfactorily fits all of the light curves with only the wavelength-dependent parameters (wavelength, limb darkening and third light) being different for each filter.
If the light curves vary in brightness throughout the cycle but the eclipse depths are quite different, then you probably have a contact or near contact system. You can again use fillout mode, setting one star's fillout to zero (0.0) and the other to a slightly less than zero value, say -0.10. Experiment with different mass ratios and different temperatures until a satisfactory fit is obtained.

If the light curve appears to be an Algol system (semi-detached with the less massive star being more evolved than the more massive star and very disparate eclipse depths), then you should probably set the program to Radii input mode. The cooler star will be at its Roche limit so set its radius to -1 (the flag that forces the star to exactly fill its inner Lagrangian surface). Estimate the mass ratio and radius of the primary star and experiment until your synthetic light curve matches your observed one.

For detached systems, again use Radii input mode and experiment with different sizes, temperatures and inclinations until you achieve success in matching your data.

6. Conclusion

Armed with the information and strategies given above, you should be able to construct a coherent model to your light curves in a very short time. With the immediate visual feedback that a light curve program like Binary Maker 3 offers you can quickly hone in to a very satisfactory model for your data. More details concerning light curve analysis, including asynchronous rotation and eccentric systems, can be found in the User Manual that comes with Binary Maker 3.

7. Acknowledgements

I want to thank the organizers of this meeting for inviting me to speak on this topic. I also want to thank Jonathan Hargis for carefully reading over the text of this paper and making very useful suggestions. My partner in crime on the Binary Maker 3 project was David P. Steelman, who translated the entire original version 2 from C into Java and greatly improved almost every aspect of the user interface.

8. References


Automated Photometry, Period Analysis and Flare-up Constraints for Selected Mira Variable Stars

Dale E. Mais
Earth Sciences Department, Palomar College, San Marcos, CA
dmais@ligand.com

Robert E. Stencel
Department of Physics and Astronomy, University of Denver
Denver, CO
rstencel@du.edu

David Richards
Aberdeen and District Astronomical Society, UK
david@richweb.f9.co.uk

Abstract
During the course of the past two years, 108 selected Mira-type program stars have been monitored to address potential flare up episodes. These include 34 M-type, 17-S type and 57 C-type Mira’s. This paper will describe the greater than 140,000 magnitude determinations that have been obtained, many closely spaced in time, which are being used to further constrain the potential occurrences of flare-up events. Random reports in the literature suggest that some Mira variables may go through flare up stages, which result in brightening on the order of several tenths of a magnitude or more, and may last hours to days in length. Very little is known about these events and their frequency, indeed, it is not clear that these events are real or instrumental phenomena. The light curves of many of the program stars show a Cepheid like bump phenomenon, usually on the ascending part of the light curve. In general, these bumps appear in longer period Mira’s (>350 days) as pointed out by Melikian in 1999. Bumps are not obvious or easily seen in visual data records, although slope changes during rising phase are seen in some cases. In order to address the reality of these events, we established an automated acquisition/analysis of a group of 108 Mira variables [M(oxygen), S and C types] in order to obtain the densest possible coverage of the periods, to better constrain the character and frequency of flare-ups. Telescope control scripts were put in place along with real time analysis. This allowed for unattended acquisition of data on every clear night, all night long, in the V, R and I photometric bands. In addition, during the course of most nights, multiple determinations are often obtained for a given star. We are grateful to the estate of William Herschel Womble for partial support of these efforts. © 2005 Society for Astronomical Science.

1. Introduction
DeLaverny et al. (1997, 1998) reported short-term brightness variations in 15 percent of the 250 Mira or Long Period Variable stars surveyed using the HIPPARCOS satellite, with the broadband 340 to 890 nm Hp filter. The abrupt variations ranged 0.2 to 1.1 magnitudes, on time-scales between 2 to 100 hours, with a preponderance found nearer Mira minimum light phases. However, the sampling frequency was extremely sparse and requires confirmation because of potentially important dust-formation physics that can be revealed.

We report here continued ground-based photometric observations of several of these objects that support and tend to confirm, the deLaverny et al. findings. According to deLaverny et al., the majority of the significant fluctuations (0.2 mag or more) occurred when the variables were at or near minimum light. Their limited sampling suggested changes on time-scales of hours and days. Initially, our observations indicated that four out of five Miras sampled (XZ Her, HO Lyr, AU Cyg and AM Cyg) were found to have significant fluctuations over these same short time-scales (at 95% confidence level and higher), based on analysis of photometric variance, and F-tests [Stencel et al. 2002].

Melikian [1999] provides a careful analysis of the light curves for 223 Miras based on Hipparcos data, finding that 82 stars [37%] show a post-minimum hump-shaped increase in brightness on the ascending branch of the light curve. Melikian advocates that differing physical processes and perhaps stellar properties – e.g. later spectral types, longer periods and higher luminosity - differentiate these stars.
Wozniak, McGowen and Vestrand [2004] reported OGLE observations that included “105,425 I band measurements of 485 Mira type variables” for an average of 217 observations per Mira spanning nearly 3 years, or roughly 1 point per Mira per 2 day interval. They did not detect any candidate events and thus set a limit of 0.038 I-band events per star per year, far lower than the ~1 event per star per year suggested by Hipparcos.

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<td>52836.6</td>
<td>1771:96 1771:3348</td>
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<td>0.003</td>
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<td>0.043</td>
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<td>CE Lyr</td>
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<td>76</td>
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<tr>
<td>HO Lyr</td>
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<td>3129:202 3129:236</td>
<td>0.041</td>
<td>0.012</td>
<td>56</td>
</tr>
<tr>
<td>V Mon</td>
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<td>4788:333 4788:509</td>
<td>0.015</td>
<td>0.045</td>
<td>61</td>
</tr>
<tr>
<td>VX Mon</td>
<td>52632.8</td>
<td>4825:861 4825:727</td>
<td>0.024</td>
<td>0.056</td>
<td>69</td>
</tr>
<tr>
<td>SW Peg</td>
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<td>1673:1014 1673:1450</td>
<td>0.014</td>
<td>0.011</td>
<td>46</td>
</tr>
<tr>
<td>V Ser</td>
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<td>939:333 939:333</td>
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<td>0.021</td>
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<tr>
<td>AM Ser</td>
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<td>1409:1014 1311:390</td>
<td>0.008</td>
<td>0.009</td>
<td>50</td>
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</table>

Table 1 Observation summary for Mira type Stars.

2. Methodology

Photometry was conducted with an Astrophysics 5.1-inch f6 refractor using an ST-10XME camera and 2x2 binned pixels and the Johnson V and R filters. Images were obtained in duplicate for each band and two reference stars used per variable star for analysis. Image reduction was carried out with CCDSOFT image reduction groups and specially written scripts for magnitude determinations, which allowed for rapid, nearly real time magnitudes to be found.

Period determination was performed using the AAVSO endorsed Peranso software suite http://users.skynet.be/fal079980/peranso/index.htm, which reports best fits using Lomb-Scargle, Bloomfield and Dworetsky methods. Initial application shows agreement with literature periods for most variables, within a few percent and with errors in fit of 1 to 20 days, depending on the data interval and phase coverage thus far. See Table 1 summary for a sample of Miras followed.

Early on it was felt that semi-automating the process was the only way to go. The use of a mount such as the Paramount along with the Suite of software by Software Bisque got the project 80% of the way there. It turns out the real consumer of time in these efforts was the magnitude determinations. This took far longer than the actual acquiring and processing of the images. Fortunately, TheSky in conjunction with CCDSOFT lends itself to scripting and a script was put together that automated the magnitude determinations. It is now not even necessary to view any of the images.

In order to further automate this process a telescope-camera-filter wheel-image reduction script was written which allows totally unattended acquisition and image reduction. In this manner, the system runs all night long cycling through target stars obtaining BVRI images of 108 target Mira variables. As a further enhancement, on the user interface window, one can specify the airmass cutoff. For example if an airmass of 2 is entered, no star on the target list will be imaged with an airmass >2. Figure 1 shows this display window for complete automated telescope control.

3. Results and Discussion

We set out several years ago to monitor a set of Miras in an attempt to “catch” a putative flare up event and obtain spectra of the event to compare to baseline spectra. As the data accumulated it became clear that these events were much rarer than previously anticipated based on some literature claims. In addition, we noted that many of these Miras had a secondary bump in their light curve that was not obvious from previous datasets. Figures 2 and 3 demonstrate this bump phenomenon. Our data set now includes over 140,000 magnitude determinations covering 108 stars over nearly a two year period, or roughly 1 point per Mira per 1 day. The most liberal interpretation of the data hints at a few, small, irregular brightenings, for an implied rate of <0.02 V,R band events/star/year.
These results might lead a sane person to abandon the experiment, but then Mighell and Roederer [2004, see poster 56.18] report flickering among red giant stars in the Ursa Minor dwarf spheroidal galaxy, including detection of low-amplitude variability in faint RGB stars on 10 minute timescales! We have also noticed this in a few stars near minimum light. An example of this is shown in Figure 4 for AM Cyg.

The data set has not been without potential flare events. Most of these events can be easily explained once the original images are examined. Since duplicate images are always obtained in each wavelength band only seconds apart, the potential "event" must be in both images to be a candidate. This eliminates most putative events. Nonetheless, a few events remain which may be real. Table 2 summarizes these small irregular events we have noticed in our data, which could not be explained by other means upon close inspection of the images. In addition Figures 5 and 6 show two examples of these putative events in the light curve of RR Boo.

Despite the challenges of verifying the existence and frequency of the "micro-flares" among selected Mira variable stars, it remains possible to speculate on causes. At least three scenarios present themselves: (1) shock induced; (2) magnetically induced, and (3) planet induced events.

(1) Many Mira variables exhibit radio maser emission arising from excited molecules of SiO in the outer atmosphere of such stars. The patchy nature of the bright SiO maser spots seen in VLBI maps varies in response to the Mira optical variation. If a consistent phase for microflares can be established, they could be related to shock propagation and interaction altitude.
Figures 2 and 3 demonstrating the presence of the bump phenomenon for RT Boo and RU Her

<table>
<thead>
<tr>
<th>Star/Period[d]</th>
<th>Event [MJD]</th>
<th>Amplitude [mag]</th>
<th>Duration</th>
<th>Type</th>
</tr>
</thead>
<tbody>
<tr>
<td>RR Boo 195 d</td>
<td>2895.621, 3105.785</td>
<td>0.22, 0.4</td>
<td>~48 hours, ~30 min.</td>
<td>V band incr at min light</td>
</tr>
<tr>
<td>RR Boo</td>
<td>&quot;summer 2001 near min light&quot;</td>
<td>~0.8</td>
<td>Not stated</td>
<td>Guenther &amp; Henson, 2001</td>
</tr>
<tr>
<td>X CrB 241 d</td>
<td>2806.682</td>
<td>0.1?</td>
<td>Days</td>
<td>V band fade, on approach to max light</td>
</tr>
<tr>
<td>X Hya 301 d</td>
<td>3105.752</td>
<td>0.3</td>
<td>days</td>
<td>V band chaos near max light</td>
</tr>
</tbody>
</table>

Table 2 Summary of possible flare events observed over a 2.5 year period
Figure 4 Presence of micro-flickering in AM Cyg between JD 3210 and 3260 in the V band?

Figure 5 Potential event #1 at JD 2895-2897

Figure 6 Potential event #2 at JD 3105-3106
(2) The same type of maser observations can be used to deduce magnetic field strengths via Zeeman line splitting. The analogy with solar magnetic phenomena [spots, flares, eruptive prominences, coronal holes and mass ejections] is compelling. However, in analogy to the R CrB phenomenon, brightness variation could also be a consequence of dust formation (fading) and dissipation (brightening) in front of a star’s visible hemisphere. Additional observations will be required to discriminate which is occurring. True flare stars include Sun-like red dwarf stars with half the surface temperature and a fraction of the solar mass. These appear to have sizeable starspots and intermittent flaring behavior. In those cases, extensive spots and concentrated fields give rise to high energy output in UV and X-rays. The latter emissions are not seen in the case of Mira variables, suggesting a strong limit to the size and strength of spots. A more “dilute” and large-scale eruptive prominence analogy might suggest measurable changes in mass loss diagnostics, such as the cores of H-alpha or Ca II K, if these could be extensively monitored for microflares (Stencel and Ionson 1979, Stencel 2000).

(3) An interesting speculation involves extending the discovery of extra-solar planets to their role around evolved stars like Mira variables. As the Mira red giant expands and engulfs its Jupiters, several kinds of accretion “fireworks” might accompany digestion (Struck et. al. 2002). However, the duration and frequency would be limited to either orbital periods or one-time events. Existing maser maps would appear to rule out large scale planetary wakes around some Mira variables, but additional observations are always merited.

4. Conclusion

Our analysis further reduces the implied rate of flare up events to <0.02 V,R band events/star/year, nearly halving the figure noted by Wozniak et.al. of 0.038 events/star/year in the I band. Analysis and some open questions which remain:

1. Precise photometry reveals bump phenomena in the light curves of many Mira variables.
2. Very few flares of a magnitude comparable to Hipparcos reporting have been seen in V, R and I bands; shorter wavelengths might hold the key to this.
3. V-R fluctuations deserve further study in these stars.
4. On what does V-R amplitude and range depend? Period? Dust type?

5. References


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Stencel, R., and Ionson, J. Stellar magnetic fields - The role of a magnetic field in the peculiar M giant, HD 4174 (1979), PASP, 91, 452


CCD Observations and Preliminary Analysis of the “End-of-the-Line” W UMa Eclipsing Binary V523 Cas

Russell M. Genet
Orion Observatory
4995 Santa Margarita Lake Road, Santa Margarita, CA 93453
russmgenet@aol.com

Thomas C. Smith
Dark Ridge Observatory
5456 Bolsa Road, Atascadero, CA 93422
tcsmith@darkridgeobservatory.org

Dirk Terrell
Southwest Research Institute
1050 Walnut Street #400, Boulder, CO 80302
terrell@boulder.swri.edu

Laurence Doyle
SETI Institute
515 North Whisman Road, Mountain View, CA 94043
ldoyle@seti.org

Abstract
V523 Cas is an “end of the line” W UMa binary; one of the most rapidly-rotating and hence magnetically active non-degenerate binaries that can exist. Determining the timescales of the resultant, rapidly changing photometric light-curve asymmetries is our primary science driver. We are also able, with considerable precision, to determine small changes in rotational periods from one season to the next. Such period changes can be caused by mass loss, the transfer of mass from one star to the other, or the Applegate effect. Third bodies can also induce changes in eclipse epoch due to the light-travel-time effect resulting from the system’s shifting barycenter. Jupiter shifts our own solar system’s barycenter some five seconds peak-to-peak, and detecting a planet orbiting V523 Cas may be within our reach. We are also looking for planetary transits across the binary pair which, unlike those across single stars, are complex and quasi-periodic. © 2005 Society for Astronomical Science.

1. End-of-the-Line W UMa Binaries

From the standpoint of stellar structure and evolutionary modeling, overcontact binary stars (Wilson, 2001) are fascinating objects, wonderful laboratories for studying structural processes that have implications far beyond understanding these binaries themselves. In them we find pairs of very dissimilar stars that are forced to coexist in physical contact. Sharing a common envelope of material, they manage to coexist with remarkable stability. Overcontact binaries are found across almost the entire range of stellar spectral types, from the hot and massive O stars to the K stars which have temperatures and masses less than the Sun.

Figure 1. W UMa schematic with Roche lobes

The most common type of overcontact binary is the W Ursae Majoris (W UMa) class where the dominant energy transport mechanism within the common envelope is convection—similar to the situation in the outer layers of the Sun. In these binaries, significant energy is transferred from the higher mass star to the lower mass star, resulting in a near
equalization of the surface temperatures, with differences usually not more than a few percent. We deduce this from the fact that, in eclipsing W UMa systems, the two eclipse depths are very nearly equal. This is a rather remarkable circumstance, given that the two stars frequently have very different masses and thus would normally have very different temperatures. Although a handful of W UMa systems have mass ratios near unity (e.g., V753 Mon at 0.97), the vast majority of them have mass ratios of 0.1 to 0.5, with the current record holder being SX Crv with a mass ratio of only 0.066. Orbital periods range from 0.22 to slightly more than one day.

We have chosen to concentrate our observations on a few of the very shortest-period W UMa systems. These include V523 Cas at 0.23 days, RW Com at 0.23 days, and CC Com at 0.22 days (the short-period record holder which we added to our program in 2005). Not only do these binaries have the shortest periods and least angular momentum of any non-degenerate binaries, they are also the coolest, as the period-color diagram in Fig. 2 reveals (Rucinski, 1993). This period-color relationship was first noted by Eggen (1961).

2. Short-Term Photometric Variability

The short-period W UMa binaries we observe are also rather unique in their high degree of magnetic activity. This is due to both their rapid rotation (dynamo effect) and their low temperature and hence deep convective envelopes with long convective turnover times (Rucinski 1993). Most late, single stars are slow rotators, and are not very active, magnetically. The late and very fast rotating W UMa binaries we are observing, on the other hand, are ideal laboratories for observing magnetic activity which can reveal itself through changes in the shapes and symmetries of the photometric light curves. Determining the degree to which these changes take place over days, weeks, and months is the primary goal of our investigation.

W UMa systems are frequently classified as either A-types or W-types. In the A-types, the more massive star is eclipsed at primary minimum while the reverse is true for the W-types. The A-types tend to be somewhat evolved, have higher total masses, and thicker common envelopes (Wilson, 1978). The W-types are unevolved objects which cannot achieve equilibrium. They are thought to undergo thermal relaxation oscillations about a state of marginal contact (Lucy, 1976; Flannery, 1976; Robertson & Eggleton, 1977). Observationally, the W-types exhibit period changes and light curve variability, as would be expected of objects that cannot achieve equilibrium. All of our observational targets are W-type W UMa systems. See the Appendix for a listing of W-type W UMa systems with periods less than 0.30 days.

Our understanding of the internal structure of over-contact binaries is surprisingly poor. Since an initial flurry of activity from the late 1960’s until the early 1980’s, very little progress has been made in understanding the structure and evolution of these stars. This lack of progress stems from both observational and theoretical difficulties. On the theoretical side, simple one-dimensional modeling (viz., Lucy, 1976; Robertson & Eggleton, 1977), necessitated by computational resource constraints, has been pushed about as far as it can go. It is clear that some very hefty three dimensional hydrodynamic (and perhaps even magnetohydrodynamic) computations are needed. This is a daunting computational task, requiring hundreds or even thousands of CPUs. Such codes are only now being developed.

On the observational side, it has long been known that W UMa systems have light curve asymmetries that are time variable. Our understanding of
the sources of these asymmetries (starspots, faculae, matter streams and, perhaps, other phenomena) and the nature of their variability are poorly understood, primarily because probing them requires large telescopes for high resolution spectroscopy or large amounts of telescope time to follow them photometrically.

While time on 2-4m class telescopes for time-intensive spectroscopy is very difficult to obtain, the task of securing high-quality photometry is much less arduous because it does not require large telescopes. Setups that can achieve millimagnitude precision photometry are even within the economic reach of serious amateur astronomers. It then becomes possible to answer the question: “What are the timescales for changes in the asymmetries of W UMa light curves?”

Answering this question is the primary science driver for our program of intensive photometry on a handful of systems over the course of an observing season. We observe a system every clear night over several months, obtaining complete light curves each night. Therefore, we can look for and adequately follow the changes in the light curves of these systems on timescales ranging from days to months. We believe that these data will prove invaluable when three dimensional models reach a maturity where they can be constrained by such observations. This time is not far off. Rather than disjointed “snapshots” of W UMa systems, we can provide “movies” of their behavior over daily, monthly and yearly timescales, enabling a more powerful feedback to the theoretical models and, ultimately, a better understanding of overcontact binary structure and evolution.

3. Intrinsic Period Changes

Three effects, intrinsic to contact binaries themselves, can cause changes in the binary’s orbital period. The first is angular momentum loss, which can be due to gravitational radiation, mass loss, or magnetic breaking. The second is mass transfer between the stars, most often from the less to more massive component. Martin and Davey (1995) suggested, based on hydrodynamic simulations, that mass might be transferred simultaneously through the neck in both directions, flowing from the hotter primary to cooler secondary in the outer portion of the connecting neck and returning, after cooling and sinking in the secondary’s atmosphere, in the inner portion of the neck.

The third is the Applegate effect. It had been noted for quite some time, that the orbital periods of a number of close binaries change directions in a nearly but not strictly periodic manner. Hall (1991) pointed out that these changes in period were correlated with changes in magnetic activity. Applegate (1991) then suggested that these changes in period modulations could be explained as “the gravitational coupling of the orbit to variations in the shape of a magnetically active star in the system.” His suggestion has, subsequently, been supported by the modeling and analysis of a sizeable group of close binaries by Lanza and Rodono (1999).

Separating the effects of real period changes from the apparent period changes can, due to the distortion of light curves by starspots, be a real challenge. Kalimeris, Rovithis-Livaniou, and Rovithis (2002) suggested how, with sufficiently densely-populated O-C curves, this distinction can be made.

4. Third-Body Induced Period Changes

It has long been known that a third star in orbit around an eclipsing binary system will produce a periodic variation in the eclipse minima timing (independent of other period-changing effects). This is caused by the movement of the binary system about the third-star / binary barycenter and the consequent light travel-time difference. That this effect could be extended to the detection of circumbinary planets was first pointed out by Schneider and Doyle (1995), and by Doyle et al. (1998) and applied to the M4-dwarf binary CM Draconis (Deeg et al. 2000). The relationship is linear, and the timing offset due to an external planet will be:

$$\Delta t = \frac{M_p a_p}{c M_*}$$

where $M_p$ is the mass of the planet, $a_p$ is the planet’s orbital semi-major axis, $M_*$ is the combined mass of the binary star system, and $c$ is the speed of light.

Through a simple propagation of errors (and assuming a triangular eclipse shape) this relationship can be reduced to observables (Doyle and Deeg 2004) and approximated by:

$$\Delta t \approx \frac{T_{ec}}{2 \Delta L \sqrt{n}}$$

where $\Delta t$ is the error in brightness measurement, $T_{ec}$ is the duration time of the eclipse, $\Delta L$ is the fractional drop in total brightness during eclipse, and $n$ is the number of observational data points taken during the eclipse duration $T_{ec}$.

As an example, this method has been applied to constrain the possible presence of a third body candidate to less than 3 Jovian masses with a period of about 1000 days around the non-contact eclipsing
binary CM Draconis (Deeg et al. 2000). The total binary mass of CM Dra is about 0.444 solar masses. In this system, eclipses last for about 80 minutes, and reach a depth (primary eclipse) of about 46%. The photometric precision was conservatively taken as 1%, and the number of data points per binary eclipse event was about 960 points (cadence of about 5 seconds). This gave a theoretical precision of about 1.7 seconds, with a rough 3-sigma detection constraining the precision in this case to about 6 seconds.

Although W Ursae Majoris systems can cause complications due to mass transfer through the first Lagrange point (the transfer of mass away from equal mass stars causing, for example, an increase in binary orbital period), this effect may not generally be expected to have a periodicity in the range of expected long-period planetary orbits. The short-period W UMa systems do have the advantage of many eclipses (the root-n term), and fractional drops might also be expected to be quite close to 50%. A disadvantage is that the duration of W UMa eclipses will be shorter, and the triangular eclipse approximation in the equation above will, in general, not be a good approximation to the shape of the eclipse minima.

5. Third-Body Transits

Planetary transits across eclipsing binary systems are particularly interesting as they produce quasi-periodic variations in the brightness of the stars, and thus are unlikely to be recognized as transit features. This is not so much because the planet is transiting the two binary stars, as the stars are orbiting “behind” the planet as it enters what one might call the “transit window.” A whole family of quasi-periodic drops in brightness can be recorded, with from one to several transit features being possible during each planet pass, depending on the orbital phase of the binaries when the planet transits. Some of these shapes are modeled in Brandmeier and Doyle (1996), including the W UMa system, XY Leo. Deeg et al. (1998) modeled shapes for CM Draconis.

Figure 3. Light curve for planetary transits at an initial binary phase of 90 degrees. One transit light curve event is slightly wider (at left) since the star and planet are traveling in the same direction across the line-of-sight. In the second transit event the star is traveling in the opposite direction from the planet.

Figure 4. Light curve for planetary transits at an initial binary phase of 0 degrees. The effect of the binary eclipse itself has been removed. Because stars are in eclipse (one stellar disc area) transit depths will be twice as deep for any given size planet at this binary phase.

Such quasi-periodic features in the light curves can be found by using a matching filter (Jenkins et al. 1996, Doyle et al. 2000). In this process—knowing the orbital phases of the eclipsing binary that are in the observational light curve—models of all possible planetary orbital phases and periods can be “matched” (in a cross-correlation sense) with the light curves to obtain a transit detection statistic of correlation. This is the optimum way to detect such quasi-periodic features, since a straightforward power spectrum will not work with quasi-periodic signals. For CM Draconis, with a period of about 1.26 days and over 1000 hours of light curve data, a correlation of over 400 million possible planet models was nec-
necessary to constrain the absence of planets of size 2.3-Earth radii or larger with periods up to 60 days with a detection probability of greater than 80% (Doyle et al, 2000)

It should be noted, as pointed out by Schneider and Doyle (1995), that planets, whose orbits are non-planar with the binary orbit but have short enough periods, will eventually have their orbital nodes precess across the line-of-sight, causing transit events that come and go. They will, that is, have transit seasons not dissimilar to the Saros cycles for lunar and solar eclipses when the Moon’s orbital nodes are aligned with the Earth’s orbital plane.

6. Observations of V523 Cas

Prior to our 2004 observations, the most recently reported V523 Cas observations were those of Samec et al (2004). Other recently reported observations include those of Zhang and Zhang (2004), Gurol et al (2003), Nelson (2001) and Lister, McDermid, and Hilditch (2000)

Samec et al (2004) described V523 Cas as having an under-luminous K0 dwarf primary and sufficiently over-luminous M2 dwarf secondary to make this a W-type system. The mass ratio, they suggested, is about 0.5 although there remains some disagreement between the mass ratios as determined by photometry and spectroscopy.

![Figure 5. Light curve for planetary transits at an initial binary phase of 170 degrees. The dotted portion of the light curve is not as deep as the solid lines due to limb-darkening; the planet does not quite transit the center of the star as it leaves the transit window (i.e., the star does not quite “catch up” to the planet).](image1)

Examples of the quasi-periodic transits produced across eclipsing binaries are shown in Figures 4 through 6. These include the two-event type of transits where one event is longer—as the star will be moving in back of the planet in the same direction—and one shorter transit event as the other star will be moving in the opposite direction from the motion of the planet (Fig. 3). Another type of transit event occurs during stellar eclipse when one star “hands off” the planet to the other star continuing the transit event (Fig. 4). This latter eclipse-transit event can last many times longer than the double transit or even multiple transits, when the first star “catches up” with the planet before it has left the transit window entirely (Fig. 5). These drops in brightness (1% for Jovian-sized planets around solar type stars) will also vary due to limb-darkening of the stellar discs and—in the case of W Ursae Majoris eclipsing binaries—with the distribution of stellar flux across the Lagrange point between the two stars as it too is transited by the planet.

![Figure 6. Dark Ridge light curve.](image2)

Some 27 complete light curves were obtained at the Dark Ridge and Orion Observatories between JD 2453253 and 2453359. (Several additional light curves were obtained prior to JD 2453253 but were discarded due to timing difficulties.) Some 14 of the 27 light curves were obtained with a 14-inch Meade LX-200GPS telescope at the Dark Ridge Observatory, while 13 light curves were obtained with a 10-inch Meade LX 200 telescope at the Orion Observa-
Both sets of observations were made with SBIG ST7XE cameras through Rc filters. Dark Ridge made 30-second integrations, while Orion, with half the glass area, made 60-second integrations. V523 Cas was observed simultaneously by both observatories on 8 nights. As somewhat more than complete orbital cycles were observed each night, a total of 32 primary minima were included within the 27 light curves.

7. Preliminary Reduction and Analysis

Our analysis, so far, has been based on very preliminary reduction procedures; hence the reduction and analysis will have to be redone once final procedures have been established. Our preliminary reduction corrected the observational times from the beginning of the integrations to their middle (a 15 second correction for Dark Ridge and 30 second for Orion), and transformed these times to heliocentric JDS.

Our photometric reduction has not yet included multiple comparison stars. Corrections for differential airmass have not been made yet, nor have the observations been transformed from our instrumental to the standard system, although observations of the secondary standards in M67 were made to facilitate such transformations.

The standard deviation of the differential C-K magnitudes, taken over the entire night at the Dark Ridge Observatory as shown in Fig. 6, was 4.9 millimagnitudes (with 30 second integrations), while the standard deviation of the best series of 20 sequential C-K differentials was 2.7 millimagnitudes. Similar values for the entire night and best series of 20 for the Orion Observatory light curve, as shown in Fig. 7, were 4.2 and 2.8 millimagnitudes respectively (with 60 second integrations). These precisions compare favorably with theoretical predictions of photometric precision (Howell, 2000).

Times of both primary and secondary minima (and maxima) were determined with Minima-23, a program kindly supplied by Robert Nelson. The effect of changing the “cutting score” (the portion of the primary eclipse “U” included in the determination) was investigated and found to change the resulting minima by a second or two, a sizeable effect from our point of view.

Thirty two times of primary minima were used to determine a seasonal ephemeris which was determined to be:

\[ \text{HJD Min I} = 2453234.6351606 + 0.23369537E \]

Calculated values (C), using the above ephemeris, were compared with observed values (O), and are shown in the O-C scatter plot below. The 1σ standard deviation of these O-C values was 4.4 seconds, while the standard error of the seasonal mean was slightly less than 0.8 seconds.

Figure 8. Seasonal O-C plot for V523 Cas.

We note that it would be inappropriate to use an already-established, multi-seasonal ephemeris to analyze our seasonal data and then, just because it showed some linear O-C slope, conclude there was a change in period within our short observing season. We could only draw this conclusion if there was a significant within-season curve to the O-C residuals which, as can be seen from the plot above, was not, at least in this instance, the case.

8. Further Reduction

To get the lay of the land, we quickly ran through our reduction and preliminary analysis as reported herein. We are now, in a more leisurely manner, refining our reduction procedures and expanding our analysis, and will report these results in subsequent papers.

One important (for us) refinement is transforming our observational times to barycentric as opposed to heliocentric time; BJD instead of HJD. The difference between these two, due to the effects of Jupiter and the other giant planets, can be as large as 5 seconds.

We are devoted to refining our approach to ensemble photometry. Ensemble photometry has been shown by Gilliland and Brown (1988, 1992), Honeycutt (1992), and Everett and Howell (2001), among others, to significantly improve photometric precision. Since we repeatedly observe the same field all night long, night-after-night, month-after-month and, we hope, year-after-year, we can afford to invest considerable effort in empirically determining, for each of our fields, the most appropriate stars to use as our comparison ensemble, and how, perhaps even under varying conditions, they might best be weighted.
There are many questions with respect to ensemble photometry we would like to investigate. Can stars of significantly different color be used to advantage if, for instance, they are the only stars of comparable brightness to the variable under study? Should all non-variable stars above some minimal brightness threshold be used in the ensemble, properly weighted to account for their noise? On long time scales, most stars are variable to some degree (Henry, 1999), but as we are primarily interested in isolating within-night light-curve changes, should we allow nightly zeropoint adjustment in our analysis and perhaps accept, as comparisons, stars which change significantly within a season but not within a night?

9. Further Analysis (Non-Parametric)

We are taking two, hopefully complimentary, approaches to our analysis. One is to fit an astrophysically-meaningful model to the observational data (see the following section), while the other is to statistically search, non-parametrically, for both regularities and changes (this section).

A portion of our analysis obtains times of primary and secondary minima as well as maxima before and after primary minima. Robert Nelson is kindly automating his Minima23 program to solve, given cutting scores and the minimum number of observational points in a determination, an entire season of light curves in one fell swoop. Such automation will allow us to explore, parametrically, the effects, for instance, of various cutting scores on the dispersion of the O-C residuals in our seasonal ephemeris. We will also be able to make a direct comparison between Kwee van Woerden (1956), Fourier low-pass, and other minima-determination techniques. With Nelson, we also plan to evaluate different approaches to estimating binary brightness (differential magnitudes) at the various times of minima and maxima.

Besides separately determining the various minima and maxima within a given light curve, we are also investigating approaches as well as treat entire light curves as entities. Various whole-curve non-parametric approaches have been utilized by Hertzsprung (1923), Tsesevich (1971), and Berdnikov (1992), and also by Turner (1999) who first suggested we consider this general approach. Our situation is somewhat unique in that we have obtained many light curves of the same system in the course of just a few months. Not expecting any significant real within-season period changes, we can treat our observations as a homogenous data set within phase space.

In our pilot season, we gathered over 10,000 points on V523 Cas, and subsequent seasons should garner many more. One analytic approach we are taking is to construct a “seasonal master” that incorporates all these points. We can then compare individual nights with this low-noise seasonal master. In another approach, we are comparing the differences between light curves, taken two nights at a time, as a function of elapsed time (orbital cycles) between the nights.

Various weighting schemes are being tried in conjunction with the above, including weights based on error estimates of individual observations and, as suggested by Ray Weymann, on the slope of the light curves in the vicinity of the individual points in question. Stetson’s (1996) suggestion with respect to alternatives to least-squares minimization is also being investigated.

10. Further Analysis (Parametric)

We intend to analyze our observations with the widely used Wilson-Devinney (WD) program (Wilson & Devinney, 1971; Wilson, 1979; Wilson, 1990). WD is a physical model rather than a geometrical one, meaning that the stellar surfaces are modeled as gravitational equipotentials rather than, say, ellipsoids. WD can model the stellar radiation fields using Kurucz atmospheres (Kurucz, 1993), simultaneously solve light and radial velocity curves, and solve for the ephemeris parameters ($HJD_0$, $P$ and $dP/dT$) without the need to resort to any time-of-minimum analysis. A version under development (Wilson, 2005) will even allow for discontinuous period changes as seen in timing diagrams of certain binaries.

For W UMa binaries of the type we are observing, typical adjusted parameters in the light curve solutions include the mean effective temperature of the secondary ($T_2$), common envelope modified surface potential ($Q_2$), orbital inclination ($i$), mass ratio ($q$), and the bandpass luminosities of star 1 ($L_1$). For partially eclipsing systems, radial velocities are necessary to determine the mass ratio, but in totally eclipsing systems the mass ratio can be determined from the photometry (Terrell & Wilson, 2005). We will model light curve asymmetries as circular spots, and determine whether there is any systematic behavior in the spots over time, i.e., can we detect any migration of the spots in longitude and/or latitude?

11. Future Observations

Last year, in 2004, most of our observations of V523 Cas were made from September 4th to 27th, with the addition of three nights in late December.
During 2005, we hope to start our observations of V523 Cas in August and continue through December. Thus we should obtain many more nights during 2005 than we did in 2004.

During 2004, our observations were made through just one filter, R (Cousins). During 2005 we plan to observe V523 Cas in two filters. As the filters are employed sequentially, the number of points obtained through any given filter per night in 2005 will only be half of what they were in 2004. However, we plan to initiate simultaneous two-color photometry, starting at the Orion Observatory as part of a summer 2005 project with Chrissy Heather, a physics major at California Polytechnic State University. Whether or not this dichroic beam-splitter, two-camera system will be ready in time for V523 Cas observations in 2005 remains to be seen.

In 2004, we focused the telescope at the start of a night’s run and rarely refocused during the night. This often resulted in significant defocus by night’s end. Although this did not appear to have a significant effect on our photometric precision (we used sufficiently large apertures in our reduction to accommodate such defocus), we hope to implement automated focus control this season with, on our brighter stars, a small degree of purposeful defocus as suggested in these proceedings by Sturm (2005).

12. Conclusions

As a class, the “end of the line” W UMa binaries are the shortest period, coolest, and most active of any of the non-degenerate binaries. As such, they offer a potentially rewarding opportunity for programs of dedicated photometry.

Our pilot program has, thus far, established that many complete light curves of these short-period systems can be gathered over an observing season in a semi-automated fashion. Preliminary and rather simplified reduction and analysis procedures have demonstrated we can routinely achieve a photometric precision of at least 3-5 millimagnitudes, and an eclipse time-of-minima determination precision of about 4 seconds per determination (1 sigma), and better than 1 second for our seasonal estimate.

As we start our second observational season, much remains to be done with our first season’s observations. We hope that ensemble reduction methods will improve our photometric precision, while the application of both parametric and nonparametric analysis techniques will reveal subtle within-season changes in the shapes and symmetries of our light curves.

We are also hopeful that refinements in our reduction and analysis will further improve the precision of our time-of-minima determinations. Although the photometric detection of a planet around V523 Cas (or any of the other short-period W UMa binaries we are observing) via either the light-time effect or transits would surprise us, we do expect to place useful limits on the presence of any such circumbinary planets.

Finally, we welcome other observers to join us in our campaign to unravel some of the mysteries of the “end of the line” W UMa contact binaries. The primary requirements are high timing accuracy (to within 0.1 seconds), good photometric precision (2-5 millimagnitudes), and compatible reduction: times transformed to barycentric JDs and differential magnitudes to the standard Johnson/Cousins system.

13. Acknowledgements

We are pleased to acknowledge the use of Robert Nelson’s Minima23 program, and we thank Eric Betz at Cuesta College for developing the Appendix. Our thanks to John Mottmann (critical discussions) and David Mitchell (suggesting we consider W UMa transits), both of California Polytechnic State University. Finally, our thanks to Brian Warner and the Society of Astronomical Science for editing and publishing this paper.

14. References


### Appendix: The Short-Period W UMa Binaries

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Extracted from Prihuta, Kreiner, and Tremko (2002).
Resurrection of a Research Telescope

Graham E. Bell
Northeast Kansas Amateur Astronomers’ League
12229 Blazingstar Rd
Maple Hill, KS 66507-8821
gebell@mindspring.com

Abstract
In 1919 Dr. Dinsmore Alter, Professor of astronomy at the University of Kansas requested funds to build a telescope to be used in the determination of asteroid and comet orbits. For years, the dream languished from lack of funding, with only incremental steps toward that goal. A telescope was built, and though not used as initially planned; it did serve to assist Clyde Tombaugh’s pursuit of a master’s degree. Eighty-six years after Dr. Altar’s request, his dream is finally coming true, but not at KU. The path from Dr. Alter’s dream to its realization provides some interesting history. © 2005 Society for Astronomical Science.

1. Introduction
For more than a century, astronomy at the University of Kansas has been following a troubled path. While education and theoretical research have done well, on campus observational research has languished.
Efforts to establish an observational program met with many more setbacks than successes.

2. Observational Facilities, the Early Years
Kansas University established an astronomy program in 1876. It was not until 1885, however, that the first observatory was built. The simple slit-roof building was constructed on land donated by the State of Kansas. Within a year of moving in a 6-inch Clark refractor and a 2-inch transit scope, Governor Robinson intervened. He rescinded the land donation leaving the Clark and transit telescopes without homes.
The astronomy department was to operate without a permanent observatory until 1919. The department did flourish in education and research under the guidance of several individuals, including Prof. E. L. Nichols. Those with a penchant for the history of physics may recognize that name, as he founded Physical Review.

3. A Decade of Struggles
Dr Dinsmore Alter was hired in 1916, but didn’t get off to a very auspicious start, taking a leave of absence before beginning his employment. He spent that time serving his country in Europe during WWI.
In 1918, upon returning to KU, one of his first acts was to request funds for a permanent observatory, a mount for the 6-inch Clark and a new 3-inch transit instrument. The total request of $3500 was funded by the state legislature. A permanent observatory was completed in the fall of 1919.
Dr. Alter had one primary field of interest, computing asteroid orbits. According to Dr. Bord [Bord 180] his goal was to create ‘an active program of asteroidal and cometary research in Lawrence.’ But such a program would require a larger telescope than the 6-inch Clark. In May of 1919, before the observatory was completed, Alter wrote a letter to Chancellor Strong, describing how he hoped to fulfill his dream. He proposed the construction of a 20-inch reflecting telescope. To sweeten the deal he offered to purchase a mirror blank, then grind and figure it. He would then donate it to KU. In return the University would purchase a mount and construct an additional observatory.
Alter acquired a blank from the Pittsburg Glass Co in 1920 and started the grinding. He set the blank aside, unfinished, when it became apparent that the University would not fund the mount for the telescope.
In 1925, Alter made another attempt, this time going to the Kansas Board of Regents. He had ammunition for this funding request. Alter obtained letters of support from some leading astronomers of the era. Among those providing such letters were W. S. Adams (Mt. Wilson), E. B. Frost (Yerkes), R. G. Aitken (Lick) and Harlow Shapley (Harvard). Alter requested $850,000. $210,000 was to be for the acquisition of a 26-inch reflector. The remainder would be placed in an endowment, used to fund hiring of two astronomers and assistants to perform the calculations. Alter failed to secure the funding, and his woes were compounded when, a year later, the observatory
was torn down to make room for an auditorium. The observatory was eventually rebuilt at a different location.

4. A Decade of Progress

In 1926 Alter met William Pitt, a retired industrialist from Kansas City. Pitt, an amateur astronomer, had made a fortune with a company he had founded. Pitt agreed to provide two large mirrors. He would acquire the blanks and grind and figure the mirrors at his own expense.

Pitt, an accomplished machinist, designed and constructed a precision grinding machine and began perfecting his polishing and figuring techniques. He converted a swimming pool in his basement into a laboratory where he and an assistant did the grinding and polishing (Fig 1). By 1928, Pitt had completed the 20-inch mirror which Alter had purchased 8 years earlier. The results were encouraging, inducing Pitt to begin a larger mirror.

Thermal stability was important, so Pitt and Alter attempted to acquire a clear quartz disk. They were unable to get the quartz, so they tried something new. They acquired a 27-inch Pyrex blank which weighed about 250 pounds and was 4.5 inches thick.

It took Pitt and his assistant more than a year to complete the mirror. In 1929 it was installed at KU. Prior to that time, Pyrex had been used for a 16-inch secondary at Mt. Wilson, but had never been used as the primary light gathering element of a telescope.

The fork mount was designed and constructed at the University. A recent KU graduate and his brother designed and machined most of the parts in the University’s machine shop. The telescope was originally installed as a prime focus instrument (Fig.2) and subsequently converted to a Newtonian (Fig 3) under the direction of N. Wyman Storer.

Even during this decade of progress, there were substantial setbacks. In 1930, just short of completion, the efforts ground to a halt. Dr. Alter wrote “a very small amount of money prevents the University of Kansas from having the prestige of the largest telescope within several hundred miles of Lawrence. When the Depression made it impossible to use more of the University’s funds, the William Pitt-University of Kansas 27-inch reflecting telescope lacked only a few minor parts: an observing chair, a complicated plate-holder and a plate measuring machine of being ready to start on its nightly career of research work.” Alter estimated that the project was halted with only $500-$800 needed for completion.

In 1935 Alter took a one year leave of absence. The following year he resigned to become Director of the Griffith Observatory. After continually promoting it for 17 years, Alter’s dream of a research observatory was still unfulfilled. It would remain so in the future.

While not fully suited for Dr. Alter’s intended research, the telescope did exist, and was functional.

About the time this telescope was first placed into service, a well known Kansas native was establishing his place in the history books. While working at Lowell Observatory in Flagstaff, Clyde Tombaugh discovered Pluto. Lacking a degree, Mr. Tombaugh returned to Kansas, where he enrolled at KU.

Dinsmore Alter had been replaced by Dr. N. Wyman Storer. One of Storer’s first acts was to have the mirror aluminized. Alter had already arranged to have this done by a KU graduate that had developed a new process. As a trial, the aluminizing was to be done without charge; KU only had to pay shipping to and from Los Angeles. Storer picked up the negotiation process, finally accepting the aluminized mirror back at the University on July 17, 1937.

Clyde Tombaugh meanwhile started work on his Masters of Mechanical Engineering. The final efforts to complete the Pitt telescope were carried out primarily by Tombaugh himself. His Master’s thesis, published in 1939 was entitled “Study of the Observational Capabilities of the University’s 27-inch Newtonian Reflector with a Program to Restore the Telescope to Pristine Condition.” In his dissertation, Clyde wrote “the writer sees little hope of much more useful research being performed until the telescope is moved to a more favorable location which would be free of obnoxious lights and dust.”

5. The Quiet Decades

Not a lot of significant use of the Pitt Telescope is found when researching observational astronomy at KU during the period from 1940 until the end of the century.

In 1944, the observatory was razed to make room for another campus building. A new one, the Tombaugh Observatory, was not completed until 1952, this time atop Lindley Hall. Lindley Hall’s Tombaugh Observatory would remain the Pitt telescope’s home until 2001.

The 27-inch Pitt Newtonian was used primarily by students for class work and by the local amateur astronomy group. It was also used as a University showpiece, giving the public the opportunity to view the night sky wonders during the frequent open house nights. Students were able to use the grating spectograph, the photoelectric photometer and an assortment of plateholders for photography. It was never used extensively as a research tool. Faculty would rely on trips to other observatories around the world for their observational research.
By 2001, the Tombaugh Observatory atop Lindley hall had fallen into disrepair. The mechanical components of the telescope needed repair. As usual, no funds were available to rebuild the observatory or repair the telescope.

6. A New Life

In 2001, our group, the Northeast Kansas Amateur Astronomers League (NEKAAL) heard of the problems with the Pitt telescope and the Tombaugh Observatory. We approached KU with the goal of obtaining the optics so that a larger telescope could be put into use for the asteroid research being conducted at NEKAAL’s Farpoint Observatory. This program had been started in 1998, primarily by Gary Hug. By 2005, the group had designations for some 340 asteroids and comet P/1999 X1 (Hug-Bell).

Thanks to the efforts of Dr. Bruce Twarog, Dr. Barbara Anthony-Twarog and Dr. Bruce Shawl NEKAAL obtained the optics on permanent loan from KU in 2001.

We (NEKAAL) now had our work cut out for us. Building a telescope around the Pitt optics called for a new design. A full Newtonian would not fit in our existing observatory and the massive mount would be expensive. Constructing a new observatory would add cost concerns.

While several of us were designing a new configuration for the Pitt optics, others were involved in
fund raising. Membership dues in our small group barely cover the annual observatory insurance costs. Everything else comes from donations. Several substantial donations were received from members, and we even got donations from KU faculty. About the time we finished raising some local funds, we completed the design of a folded f5.5 Newtonian. Our funds were insufficient!

Several grant applications provided no funds, including an application to NASA. Those who reviewed the NASA grant application encouraged us to apply again, after we made some changes. We were to find another person to build the scope and needed to develop a better implementation plan.

A search and queries of others involved in minor planet research led us to Jerry Foote of ScopeCraft. We NEKAAL and ScopeCraft) put together a timeline for construction and developed a contract with ScopeCraft. Both became a part of the next NASA grant application. In February of 2004, we received word that the grant was being funded, approximately three times as much money as we had requested in our first NASA grant application. This was initially met with cheers from the NEKAAL members… ‘We got the money.’ Shortly the attitude was one of ‘Oh, no. Now we have to perform and accomplish all we promised in the grant application.’

In July, the funds were finally available and construction started.

Clyde Tombaugh was a Kansas native, the telescope would be installed around the time of the 75th anniversary of his discovery of Pluto. And he had a major role in refurbishing this telescope the first time. The new scope would therefore be named the Tombaugh Telescope.

With a newly aluminized primary, and an SBIG 1k x 1k CCD camera (STL1001e), the installation of the Tombaugh Telescope was started at Farpoint Observatory on March 10, 2005 (fig. 4). The first NEO observations using this scope were submitted to the Minor Planet Center the night of March 15-16 2005, some 86 years after Dinsmore Alter started his efforts to build a minor planet research facility.

As initially configured, the new Tombaugh telescope failed to live up to expectations. The design had been changed from one with only a primary and a tilted secondary to one which used a tertiary. That reduced the effective aperture somewhat, as the tertiary blocked some of the light path. Baffling was not sufficient resulting in rather poor images. This was compounded by our failure to obtain adequate flat frames during the rush to become productive. With the internal reflections our traditional method of obtaining flats with a halogen light reflected from a screen was not going to work. Night sky flats appear to be the only practical way to assure good flats. But even those would be inadequate if internal reflections vary with telescope pointing.

As this is being written, ScopeCraft is building parts which will allow us to remove the tertiary, placing the camera within the OTA. This will increase the effective aperture. ScopeCraft is also reworking the baffling.

From figure 4, it is easy to see why a folded design was required. The scope almost fills the observatory. Note also the small clearance between the OTA and the roll off roof. With the changes currently planned, the scope will have an effective aperture of approximately 24 inches, which should substantially surpass our old 12-inch SCT capabilities.

NEKAAL also obtained an E/PO grant from NASA, with which we purchased two scopes (8 and 12 inch GPS models) to be used for our outreach activities. The 14-inch will be housed in a new, smaller observatory currently being designed. The new observatory is funded in part by the NASA grant.

A year ago NEKAAL owned no telescopes. All the scopes in use belonged to individual members. Now, finally, NEKAAL owns three scopes.

The 27-inch Tombaugh telescope being installed at NEKAAL’s Farpoint Observatory.

7. Conclusion

It took almost 90 years, but Dinsmore Alter’s dream of a facility for minor planet research has become a reality. Not at KU as he envisioned, but at a location within 70 miles of KU, and at a dark site as recommended by Clyde Tombaugh.

NEKAAL’s dream of owning a large (for us) research telescope has been fulfilled.

Some of us became pretty proficient at grant writing.
8. Acknowledgements

Most of the information used in this paper, especially for the time period through 1939, is from the paper by D. J. Bord [Bord 1980] which was published when he was living in Kansas. Dr. Bord, now with the University of Michigan, graciously let me draw heavily on that paper without restriction.

The Tombaugh telescope described in this paper was funded in part by a $56,060 NASA grant. Construction of that telescope was done by ScopeCraft, Kanab, Utah. The additional observatory soon to be built and the additional telescopes recently acquired were funded in part by the NASA E/PO grant. Substantial contributions were received from the Marvin Kessler estate and from Mrs. Ruth Fink of Topeka, KS.

9. References

Asynchronous Binary Asteroids - Pravec

Photometric Survey for Asynchronous Binary Asteroids

P. Pravec
Astronomical Institute AS CR, Fricova 1, CZ-25165 Ondřejov, Czech Republic

Abstract

Asynchronous binary asteroids have been found to be abundant among fast-spinning near-Earth asteroids (NEAs) smaller than 2 km in diameter; Pravec et al. (2005, Icarus, submitted) derived that 15 +/- 4 % of NEAs in the size range 0.3 to 2 km are binary with the secondary-to-primary mean diameter ratio >=0.18. The early results from the surveys of the Vesta family and the Hungaria group (Ryan et al., 2004, Planet. Space Sci. 42, 1093; 2004, Bull. Amer. Astron. Society 36, 1181; Warner et al., 2005, IAU Circ. 8511) suggest that the population extends beyond the region of terrestrial planets, but with characteristics shifted to larger sizes and longer periods; the four known binaries in the Vesta family/Hungaria group are 3 to 6 km large and they have primary rotation periods in a range of 3 to ~4 h, i.e., on the tail of the distribution of primary rotation periods of NEAs. The comparison suggests that formation and evolution mechanisms of asynchronous NEA and main-belt binaries may be similar and are related to their fast spins and rubble-pile structure. None of the current theories of their formation and evolution, however, explains the observed properties of both NEA and main-belt asynchronous binaries in full. We have established a collaborative observational program, called "Photometric Survey for Asynchronous Binary Asteroids" to discover and describe asynchronous binaries over a range of heliocentric distances from NEAs through Mars-crossers to inner main-belt asteroids. One new binary Amor asteroid, 2005 AB has been found during the first few months of the survey operation (Reddy et al., 2005, IAU Circ. 8483), and we have obtained follow-up data for two other binary systems. I outline the motivations, the technique, and the strategy of the Survey.

1. Introduction

Binary systems among small, especially Near-Earth Asteroids (NEAs) have been observed since mid-1990's by lightcurve photometry technique and since 2000 by radar (see Merline et al., 2002, and references therein). Most binary near-Earth asteroids are asynchronous systems; their primaries rotate with periods shorter than the mutual orbital periods. This property played a key role in establishing the efficient technique of their detection with lightcurve observations (Pravec et al., 2005b, and references therein, also outlined in Section 2). The up-to-date list of known binary NEAs is available on

http://www.asu.cas.cz/~asteroid/binneas.htm

A comparison of properties of the binary NEA population with those of binaries observed in other groups of asteroids lying outside the region of terrestrial planets is needed to understand relations between the different populations of binaries and to provide further data for constraining theories of their formation and evolution. Pravec et al. (2005b) have done some initial work on this. They compared the data on NEA binaries with data obtained by Ryan et al. (2004a, b) and Warner et al. (2005; personal communication) for two binary Vestoids and two binary Hungaria asteroids, respectively, and obtained a few interesting hints and constraints that suggest that the Vesta family/Hungaria group binaries (henceforth called "VH binaries") were formed and evolved by same or similar mechanisms as some or all binary NEAs.

Theories of formation and evolutionary processes of binary asteroids were summarized in Merline et al. (2002). None of the processes, however, explains the observed properties and abundance of binary NEAs as well as the VH binaries fully. Bottke and Melosh (1996a, b), Richardson et al. (1998), and Walsh and Richardson (2004) examined the tidal effect of planetary encounters on gravitationally bound aggregates and they proposed that Near-Earth Asteroids of a rubble pile structure (with zero global tensile strength) can evolve into co-orbiting binaries. While this mechanism may be responsible for some NEA binaries, it does not work in asteroid groups lying beyond the region of terrestrial planets.

Of the other two mechanisms mentioned in Weidenschilling et al. (1989) and Merline et al. (2002), neither seems to fit with the observed properties and
abundance of binary NEAs as well as VH asteroids. The cratering ejecta mechanism predicts irregularly shaped, elongated primaries. The disruptive capture mechanism predicts no initial preference for rapid rotation of primaries and no correlation with the primary's shape.

Recently it was proposed that binary systems with properties of the NEA as well as the VH binaries might be created by rotational fission of small rubble pile asteroids that were spun up by the YORP effect (Rubincam, 2000; Bottke et al., 2002).

To obtain a thorough understanding of the population of asynchronous binary asteroids over the range of heliocentric orbits and to constrain theories of their formation and evolution, we have established the project "Photometric Survey for Asynchronous Binary Asteroids".

2. Lightcurve technique for asynchronous binary detection

The technique for binary asteroid detection was reviewed in Pravec et al. (2005b). We outline its principles below.

An asynchronous binary asteroid is an object of two asteroidal bodies in a mutual orbit with at least one of them rotating with a period different from their orbital period. Photometric observations of such system can reveal signals with the two (or more) different periods. Each of the two bodies scatters sunlight that produces its own rotational lightcurve, and mutual events occur in favorable geometric conditions when the Earth or Sun is close enough to the mutual orbital plane of the system. The two rotational lightcurves add linearly into a combined lightcurve that can be represented as a linear addition of two Fourier series

\[ F(t) = F_1(t) + F_2(t) \]

\[ F_1(t) = C_1 + \sum_{k=1}^{m_1} \left( C_{1k} \cos \frac{2\pi k}{P_1} (t - t_0) + S_{1k} \sin \frac{2\pi k}{P_1} (t - t_0) \right) \]

\[ F_2(t) = C_2 + \sum_{k=1}^{m_2} \left( C_{2k} \cos \frac{2\pi k}{P_2} (t - t_0) + S_{2k} \sin \frac{2\pi k}{P_2} (t - t_0) \right) \]

where

- \( F(t) \) is the total reduced light flux at time \( t \)
- \( F_j(t) \) are the reduced light fluxes of the components at time \( t \)
- \( C_j \) are the mean reduced light fluxes of the components
- \( C_{jk}, S_{jk} \) are the Fourier coefficients
- \( P_j \) are the rotation lightcurve periods
- \( t_0 \) is the zero-point time (epoch)
- \( m_j \) are the maximum significant orders

(see also Pravec et al., 2000, and references therein).

The two constant terms are added to \( C_0 = C_1 + C_2 \) which is fitted in analysis. Note that using the representations with the above formulas, we assume a principal axis rotation for each of the components; a non-principal axis rotation would produce a complex lightcurve (see Pravec et al., 2005a).

Mutual events produce attenuations that are superposed to the combined rotational lightcurve of the two system's components. A shape of an individual attenuation event depends on instantaneous orientations, shapes, and surface brightness distributions of the two components as well as on illumination and viewing geometries of the system at the time of the event. Total events have, however, a few characteristic features:

- A plateau of constant brightness attenuation is seen during the total secondary event after the primary variation is subtracted from the lightcurve data.
- Slopes of increasing and decreasing branches (occurring during orbital phases where the bodies partially obscure one other) are steeper than slopes of the rotational lightcurve of the secondary. It is due to the fact that during the partial phases prior and after the total events, the rate of obscuration of the occulted/eclipsed body (in units of area per period) is greater than the rate at which parts of the secondary rotate into/out of view.

The features allow resolving between a rotational component feature (minimum) and a mutual event even in a binary system with the secondary rotating synchronously with the orbital motion. Onsets/offsets of the partial phases of the total events produce fast, large changes of the slope of the combined secondary's lightcurve and the mutual event attenuations.

The depth of the attenuation during the total secondary event provides an estimate of the constant term \( C_0 \), which leads to resolving the degeneration of the two constant terms in the fitted term \( C_0 \).

For same albedos and phase effects of the two bodies, the depth of the total secondary attenuation is
related with the ratio of their mean projected diameters with the formula

$$\frac{C_2}{C_0} = \left[ \left( \frac{D_a}{D_b} \right)^2 + 1 \right]^{-1}$$

When only partial events occur, we do not see a plateau of constant brightness attenuation in the secondary event. Also, the slopes of increasing and decreasing branches of the events may not be distinctively greater than slopes of a rotational component of the secondary, making an immediate resolution between secondary's rotational component's minima and mutual events less certain in some cases. In such instances, further observations made in changing geometric conditions of the system with respect to Earth and Sun should bring an answer when the asteroid moves into a more favorable geometry that produces deeper events.

In a case where we see a two-periodic lightcurve that is well described with the additive two-period Fourier series but no clear mutual attenuation events (i.e., either the mutual events do not occur in the given geometric conditions, or they are too shallow so that we cannot distinguish them from the secondary's rotational lightcurve minima), we consider the asteroid as a probable binary system as well.

The possibility that it could be a tumbling asteroid has to be investigated but unless a significant signal in linear combinations of the two main frequencies is found, the lightcurve consisting of the two additive components favors the binary interpretation (see Pravec et al., 2005b).

An example of lightcurve data of asynchronous binary system is presented in Fig. 1 and 2 (reprinted from Pravec et al., 2005b). The two figures show the observational data obtained for (65803) Didymos during two different intervals. Both figures present the data in three different forms:

- The (a) part of each figure shows the original data folded with the orbital period, reduced to unit geo- and heliocentric distances and to a given phase angle using the H-G relation with the best fit value of G.
- The (b) part presents the data with the primary variation component removed; \((F_1(t)-C_1)\) was subtracted from each data point. The displayed data therefore represent the secondary lightcurve component with superposed mutual attenuation events, with the mean primary light flux \(C_1\) present as well. (The full fitted constant term \(C_0=C_1+C_2\) was left there and not subtracted from the data plotted in the figures.).

- The (c) part shows the primary rotation component; it presents data taken outside the mutual attenuation events with the secondary variation \((F_2(t)-C_2)\) subtracted and folded with the primary period.

We point out that while all the fits and subtractions of components were done in linear, flux units, the figures are plotted in magnitudes in order to present the data in the standard units.

Fig. 1: Lightcurve data of (65803) Didymos of 2003-11-20.9 to 24.2 folded with the periods of 11.91 h (a, b) and 2.2592 h (c), and \(G=0.20\). (a) The original data showing both lightcurve components. The additive Fourier series with the periods of 2.2592 h and 11.91 h was fitted to the primary and the secondary rotation data. (b) The long-period component showing the mutual events and the secondary rotation lightcurve. The primary lightcurve component was subtracted. (c) The primary lightcurve component. The epoch of the primary component’s plot is the same as the epoch of the long period component’s plot (all times are JD [UTC] light-time corrected).

Fig. 2: Lightcurves of (65803) Didymos of 2003-11-26.2 to 12-04.1. See caption to Fig. 1. The best-fit synodic primary period was 2.2593 h.

3. Properties of the Population of Binary NEAs

Pravec et al. (2005b) analyzed observational selection effects of their survey for binary NEAs and
Asynchronous Binary Asteroids - Pravec

derived characteristic properties of their population. A few figures from the Pravec et al. (2005b) are reprinted below.

I summarize their conclusions in following.

- Binary systems with \( D_s/D_p > 0.18 \) concentrate among NEAs smaller than 2 km in diameter; the abundance of binaries decreases among larger NEAs. See Fig. 3.

- Secondaries show an apparent upper size limit of \( D_s = 0.5 \text{-} 1 \) km. Systems with the secondary-to-primary mean diameter ratios \( D_s/D_p \leq 0.5 \) are abundant while larger satellites are less frequent. See Figs. 3 and 4.

- Primaries have spheroidal shapes and they rotate fast, concentrating in the range of periods 2.2-2.8 h and with the tail of the distribution in the range 2.8 to ~4 h. The fast rotators are close to the critical spin for rubble piles with bulk densities about 2 g/cm\(^3\). See Fig. 5 and 6.

- Orbital periods show a cut-off at \( P_{orb} \sim 11 \) h; closer systems with shorter orbital periods are rare or non-existent, which is apparently consistent with the Roche’s limit for strengthless satellites. See Fig. 4.

- On average, secondary shapes are more elongated than primaries. Their rotations appear to be mostly synchronized with the orbital motion in close systems with \( P_{orb} < 20 \) h, but it appears that some systems with larger separations have unsynchronous secondary rotations.

- The available data do not provide evidence on whether the asynchronous binary population remains the same or changes, in abundance or in formation mechanism(s), with perihelion distance beyond \( q = 1.05 \) AU. A comparison with the four binaries known in the Vesta family and the Hungaria group suggests that the population extends beyond the region of terrestrial planets, but with characteristics shifted a bit to larger sizes and greater periods.

The characteristics of the binary NEA population indicate a formation mechanism closely related to their fast spins and rubble pile structure. However, neither the observational data nor predictions from the theories are detailed or thorough enough to distinguish whether the NEA binaries were created by a mechanism related to their near-Earth orbits (e.g., by tidal splitting during planetary encounters) or by some other mechanism that also works in more distant orbits (e.g., the spin-up YORP effect).

Fig. 3: Secondary-to-primary mean-diameter ratio vs. primary diameter for the 12 regularly detected binary NEAs (filled circles) within the photometric survey, and for 10 additional systems detected primarily or exclusively by radar (see Pravec et al., 2005b). Secondaries concentrate at and below \( D_s = 0.5 \) km.

Fig. 4: Secondary-to-primary mean-diameter ratio vs orbital period for the 12 regularly detected binary NEAs (filled circles) within the survey by Pravec et al. (2005b). Three other systems with \( D_p > 0.3 \) km are included, \((66391)\ 1999\ KW_4\), \((69230)\ Hermes\) and \(2002\ CE_26\). They were measured mostly with radar. The horizontal dashed lines at \( D_s/D_p = 0.15 \) and 0.5, respectively, in this as well as in Fig. 3, indicate the photometric detection lower limit and the apparent upper limit of the range of concentration of the secondaries. The vertical dashed line at \( P_{orb} = 11 \) h indicates the apparent cut-off in the orbital periods.
Asynchronous Binary Asteroids - Pravec

4. Photometric Survey for Asynchronous Binary Asteroids

Since December 2004 we have run an extended survey looking for binary asteroids. A group of interested and devoted asteroid photometrists that work in a coordinated way and use a strategy as outlined below was established. They coordinate their observations internally; some more general information has been placed on the www pages

http://www.asu.cas.cz/~asteroid/binastphotosurvey.htm

The observational strategy used in this Survey is an enhanced version of the strategy that Pravec et al. (2005b) used for detection of NEA binaries during 1994-2004 and that allowed them to model selection effects of their survey.

A central point of the strategy is to cover a targeted asteroid thoroughly on a few nearby nights so that its (primary) period can be estimated uniquely and a potential attenuation feature, or a secondary period, resolved immediately. A fast reduction and analysis of the observations, basically before beginning of following night, is a necessary condition to achieve the goal.

When an attenuation feature or a second period is found in the data, a few other stations participating in the Survey are asked to collaborate on taking further data needed to describe the binary system.

4.1 Targeted asteroids

While binary NEAs concentrate in the size range below 2 km and their abundance decreases significantly above 2 km (though a small fraction of them may be as large as ~4 km), the VH binaries appear to have the apparent upper limit shifted to larger sizes. The four known VH binaries are 3 to 6 km in size.
We therefore extended the Survey to asteroids with sizes up to 10 km. Since an actual size and albedo of such small asteroids are usually unknown, we have to estimate the size from the absolute magnitude using an albedo value assumed according to a typical albedo in the given group of asteroids. For example, most asteroids in the Vesta family as well as in the Hungaria group of asteroids have a high albedo of \( p_V = 0.3 \) to 0.4, so the size limit of 10 km converts to a limit in \( H \) of ~11.8. In other groups of asteroids (NEAs, Mars-crossers, other inner main-belt asteroids), a typical albedo is about \( p_V = 0.2 \), which converts the 10-km size limit to \( H \approx 12.4 \). Given the uncertainties of a few tenths of magnitude in the majority of available \( H \) estimates, we use a practical limit for selecting targets for the Survey of \( H > 12 \).

The selection procedure takes into account observational conditions of the asteroid for the given station during a few weeks. Observational windows lasting for a week or less are too narrow, as it might be difficult to get sufficient follow-up for the asteroid, if discovered to be binary, in the narrow window.

During the observational window, the asteroid should be observable at airmass lower than 2 for at least a few hours on each night. The brightness and motion of the asteroid should allow the observer's
system to obtain photometric errors of 0.03 mag or less during the observational window.

Ideally, the selection procedure should not consider previous lightcurve observations of the given asteroid. In practice, it suffices to check whether the asteroid has been covered thoroughly on at least one past apparition. If it was not, the asteroid might be targeted within the Survey again so that a sufficient coverage is obtained during the new apparition.

4.2 Time coverage

Generally, long nightly runs are much more preferred than short ones. This is due to the fact that the orbital periods are all relatively long (with the lower limit of ~11 hours but some have orbital periods in a range of tens of hours), so long nightly runs increase the probability of catching a mutual event. Shorter runs may be usable as well, but runs with durations below 2 hours are of little use.

A minimum requirement on the coverage of primary rotation period is that each rotation phase has to be covered twice. Gaps in the coverage of primary period shorter than ~0.5 hour can be tolerated as mutual events of significant depths last typically for 1 or 2 hours, but any longer gap needs to be covered with further data on another night. Since the required minimal double coverage of the primary rotation period would mean an extensive length of observations for longer periods, and considering that the known asynchronous binary NEAs and VH asteroids have $P_{\text{prim}} < 5$ h, we established a practical upper limit for the double coverage to be applied only to asteroids with rotation periods shorter than 10 hours.

So, the strategy actually is: 1) establish a unique period solution for the targeted asteroid, 2) complete the double coverage of the derived rotation period if it is less than 10 hours.

A minimum number of nights that will be needed for the particular asteroid therefore cannot be well planned until data on the first night, sometimes several, are obtained. The requirement to obtain a unique solution of the rotation period is sometimes fulfilled in a single night (if it is long enough so that more than one cycle is observed), but more typically it takes 2 or 3 nights. The requirement to cover each primary period phase twice means that the total length is always at least $2P_{\text{prim}}$, but it usually takes about $3P_{\text{prim}}$ due to overlaps and an “interference” between $P_{\text{prim}}$ and one day. So, the station that targets the particular asteroid has to plan to 20 to 30 hours total coverage during at least 3 nights, though it may turn out that the object is finished sooner if its rotation period is significantly shorter that 10 hours.

If the station participating in the Survey cannot complete the coverage of the targeted asteroid for any reason, or if the station gets less than two full observing nights, it may still contribute usefully to the Survey. This can be either by collaborating on an asteroid observed from another station so that the unique solution and/or the double coverage of the asteroid’s rotation period is obtained faster, or by targeting a binary asteroid discovered earlier for which further follow-up observations are needed.

4.3 Calibrated vs. Uncalibrated data

Data calibrated on the standard system (Cousins R band is preferred) or mutually linked to a level of 0.02 mag or better are most useful. Relative (differential) measurements are useful as well if they cover features that allow solving for the magnitude zero point. In practice, it turns out that relative runs longer than 4 hours are almost always usable while runs shorter than 2 hours are rarely useful.

5. Concluding remarks

The estimated abundance of asynchronous binaries among NEAs of 15 +/- 4% and their estimated detection probability of the Survey of about 0.40 (Pravec et al., 2005b) means that about 6% of asteroids targeted within the Survey will actually be resolved as binary. The early data for VH binaries suggest that the abundance of binaries among them is similar to that of NEAs. So, we may expect to detect a binary in 1 of ~20 asteroids targeted within the Survey. Since we run the observations in a way that allows modeling the selection effects of the Survey, non-detections are as important as binary detections.

After we cover a few hundred targets during the Survey and discover more than 10 new binary systems over a range of heliocentric distances from NEAs to inner main-belt asteroids, it will be the right time to do a thorough analysis of the selection effects of the Survey. This will be done using methods similar to those used by Pravec et al. (2005b) for their survey of NEA binaries, yielding an estimate of the properties of the population of asynchronous binary asteroids over the range of heliocentric orbits.

In the meantime, publications for individual resolved binaries as well as interesting non-detections may be done as found suitable by the principal observer/station for each individual case. A summary of the results for each targeted asteroid can be placed on dedicated web pages after finishing the target within the Survey. And, of course, each discovered binary is published on the IAU Circular immediately after the discovery.
6. Acknowledgements

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7. References


Asteroid Phase Curves: New Opportunities for Amateur Observers

Richard Miles
British Astronomical Association
Grange Cottage, Golden Hill, Stourton Caundle, Dorset DT10 2JP, United Kingdom
rmiles.btee@btinternet.com

Abstract
The characteristics and sparsity of accurate asteroid phase curve data are highlighted. Though a challenging proposition for CCD photometricists, the author puts forward an observing methodology exploiting the Hipparcos Catalog as a source of homogeneous V magnitudes to perform 'quasi-absolute' photometry to high accuracy. A procedure for transforming unfiltered photometry to V magnitudes is described. Application of the quasi-absolute photometry methodology to variable star measurement is demonstrated. © 2005 Society for Astronomical Science.

1. Introduction

Much valuable work is now being carried out by amateur observers to characterize the rotational rates of asteroids, and the number of asteroids which have been studied by all observers has now surpassed the 2000 mark with reasonably well-known periods available for some 1658 objects (reliability code 2 or higher according to Harris (2005)).

However, one area that has taken a back-seat over recent years has been that of asteroid phase curves and in particular observation of the opposition effect at low phase angles. So far, only about 30-40 asteroids have been studied down to very low phase angles (0.3 deg or less), which contrasts strongly with the number investigated from the pure rotational standpoint. One explanation for this has been the move to using small-area CCD cameras, by which both the asteroid and comparison star(s) can be imaged at the same time, facilitating differential photometry. Although it may happen that an accurate reference star can be found in the same CCD field as an asteroid on one particular night, invariably the asteroid will move such that on other nights the chances are that no such star will be available. Those stars that are used as comparisons tend to be relatively faint and as such do not have accurate catalog magnitudes (say to better than ±0.05 mag). The consequence of this observing methodology is that it is not possible to construct the phase curve since individual nights’ observations can not be put on a standard magnitude footing.

One major consequence of the shortage of phase curve data is that theoretical studies of the nature of the opposition effect are currently held back. Space-probe data and ground-based observations of other Solar System objects are used to a limited extent. However, better coverage of the asteroids, especially of the brighter ones, at very low phase angle would help to provide further insight into the light-scattering properties of surface regoliths on airless worlds amongst the various families of asteroids.

The present paper presents a novel approach to absolute photometry as exemplified by work on variable stars, and aims to encourage would-be observers to take up the challenge of adopting an absolute V-band photometric approach to their studies of asteroids.

2. Photometric Reference Stars

Before proceeding with a discussion of this topic, I need to point out one unfortunate fact, which distinguishes the lot of the amateur compared with that of those making observations from traditional mountain-top observatories staffed by professionals. The amateur has to settle for whatever skies are available from his/her backyard observatory. Yes, occasionally a truly clear night comes along from time to time, when it is possible and practical to use standard stars such as those of Landolt and Cousins. The amateur has to settle for whatever skies are available from his/her backyard observatory. Yes, occasionally a truly clear night comes along from time to time, when it is possible and practical to use standard stars such as those of Landolt and Cousins. But my experience in the UK and Europe is that such nights are relatively rare, arriving about 20-30% of the time. If you want to follow an asteroid through its apparition, it is therefore not practicable to use the well-tried and trusted method for absolute photometry in which it is assumed that the atmospheric extinction remains relatively constant over periods of say one hour or so. If you did wait for only truly clear, 'photometric' nights, then any phase curve you managed would likely be very incomplete. Note that asteroids are at low phase angle for only a week or
so, when you cannot afford the luxury of picking and choosing which nights to observe.

So what options are available in terms of photometric reference stars when observing moving objects such as asteroids? Most CCD fields available to amateurs using medium-sized instruments (15-40 cm aperture) span about 10-30 arcmin and as such typically encompass a few Tycho stars, and a good number of UCAC-2 stars. However they are unlikely to contain either Landolt, Cousins or other accurately measured stars as listed in the General Catalogue of Photometric Data, Mermilliod et al. (2000). Although the astrometric accuracy of the UCAC-2 catalogue is second to none, its photometric accuracy when transformed to V magnitude is only good to about 0.2 mag with systematic bias being the main problem. Tycho-1 photometry was also marginal. Figure 1 illustrates the difference between Landolt V magnitudes and Tycho-1 Vt magnitudes for stars brighter than V=10, the scatter being close to 0.05 mag.

Figure 1. Difference between Landolt V and Tycho-1 Vt magnitudes vs. B-V color.

Figure 2. Correlation between Landolt V and Tycho-2 Vt magnitudes vs. B-V color.

Tycho-2 is better, provided that the stars used are relatively bright (see Figure 2). For stars of V<10, the photometric accuracy is potentially about 0.025 mag (st. dev.) for V magnitudes derived from Tycho Vt and Tycho Bt data. For fainter stars useful as direct comparison stars, accuracies of 0.2 mag or worse are commonplace. It should be pointed out, however, that the Tycho data on more than a million stars were obtained using the star mapper instrument on board the Hipparcos spacecraft. The primary Hipparcos instrument yielded superior photometric (and astrometric) precision for a subset of some 118,000 stars.

Indeed, one advantage of asteroid photometry compared to variable star work concerns the limited range of color of asteroids, most having a B-V in the range 0.6-1.0. So if comparison stars are selected from Hipparcos stars only having colors in the range 0.2<B-V<1.0 then the standard deviation about a second-order polynomial fit reduces to a mere 0.004 mag and opens up the possibility of their use as secondary photometric standards.

Bessell (2000) has extracted Hipparcos and Tycho magnitudes for several hundred of the Cousins E-region standard stars and carried out a similar correlation fit of Hp magnitudes against Johnson V magnitude. Interestingly, the mean fit listed in Table 1 of Bessell’s paper agrees with that shown in my Figure.
3 above to one or two millimagnitudes over the range 0.0<B-V<1.0.

3. **Absolute Photometry based on V Magnitudes derived from Hipparcos: An independent test of precision and accuracy**

Although it is possible to model say Hipparcos data against Landolt or Cousins standards and achieve an internal precision of a few millimagnitudes, the questions for photometricists are, “What is the accuracy of Hipparcos data for *individual stars*?” and “What is the measurement accuracy in practice using amateur equipment?” To test this out, I have examined results obtained from my observatory on extinction and exo-atmosphere zeropoint determinations for all nights where extinction was relatively low (kv<0.28) and where at least 4 Hipparcos stars were measured at various airmasses.

Table 1 summarizes the resultant data which is based on measurements on 16 nights during the most recent observing season using a Takahashi FS60C refractor (60 mm aperture) equipped with V-filter and Starlight Xpress SXV-H9 CCD camera. The stars used comprised 60 different Hipparcos stars selected from a subset of 311 stars ranging in brightness from about V=7 to V=9 and having a B-V color in the narrow range, 0.80-0.85. Measurements were based on 10 images of each star. Plots of V-v versus airmass were linear up to airmasses of about 5. Extrapolating the linear fit to zero airmass yielded the exo-atmosphere zeropoint, Zv (for stars of this color) which on a night-to-night basis exhibits a standard deviation of about 0.01 mag. Indeed, if only results where 6 or more stars were used for each extinction plot, (8 plots in all) the scatter in Zv is a mere 0.007 mag, which is strong evidence that the Hipparcos star V magnitudes are accurate to a high order. The zero-point values were also validated on the night of 2004 Sep 18, when for example 8 Landolt stars in the selected area SA-111 were used to derive exo-atmosphere zeropoints, the V-band results obtained by the two methods differed by 0.005±0.007 mag. Also listed in Table 1 is the mean deviation of measures on *individual* stars for a total of 18 extinction plots expressed as the average residual from the linear regression fit. The average mean deviation based on 60 Hipparcos stars was found to be 0.008 mag. This is a very robust independent test of the accuracy of V magnitudes derived from Hipparcos data and supports the findings described above. Note that a number of observational and instrumental factors such as varying sky transparency, photon statistics, drift in electronics, variation in transmission of the optics, all conspire to degrade measurement accuracy, so the underlying accuracy of the source data must be of the order of 0.005 mag or better for stars similar in color to asteroids. Given that the Hipparcos dataset contains several tens of thousands of stars suitable for use as comparison stars, it represents the best source of reference stars to be exploited in performing absolute photometry of asteroids.

4. **Observing Methodologies: Quasi-absolute Photometry**

It is not that obvious how best to exploit Hipparcos as a source of accurate reference stars for V photometry of asteroids. The main limitations are; (a) Hipparcos stars are relatively bright and would therefore tend to saturate in CCD images unless very short exposures are used, and (b) although there are on average 1 or 2 suitable Hipparcos stars per square degree of sky, most CCD fields are small and will not usually include such a star.

We also need to consider sky conditions since these will also affect the choice of optimal observing methodology. Provided skies are of photometric quality, it is always possible to offset the telescope, defocus slightly and utilize say 20-30 short exposures to image a suitable Hipparcos star for the purpose. However, most skies are not photometric and so preclude this approach. A pragmatic solution to this dilemma is to use a secondary telescope of smaller aperture and wider field of view to calibrate against.

In my own case, I opted for a 60-mm aperture fluorite refractor having a focal length of 355 mm for use as a secondary scope. Images with this small telescope are taken using the same CCD as for my main 280-mm Schmidt-Cassegrain but encompass a wide field of about 1.5 square degrees of sky, ideal for capturing a few Hipparcos stars as photometric comparisons. The small aperture also helps when it comes to avoiding saturation when the relatively bright Hipparcos stars are imaged. A further help in this respect is to operate the ancillary scope defocused slightly so as to reduce the maximum pixel value by a factor of 3 or so relative to the perfectly focused stellar image – this ‘trick’ effectively extends the useable dynamic range by a magnitude or two. Of course, I equipped the smaller scope with a high-quality V filter and my flatfields are typically based on a large number (100 or so) sky flats taken at dawn.
There are several observing methodologies that can be used with such a system. If the asteroid is relatively bright, say having V<13 then it is possible to image the asteroid and Hipparcos star with the 60-mm aperture V-telescope directly and perform an analysis of the images taking into account atmospheric extinction and transformation to the standard Johnson V magnitude system. This is a form of ‘quasi-absolute’ photometry, ‘quasi’ in as much as traditional standard stars of the Landolt or Cousins variety are not used and it is not necessary to shift the telescope to image the Hipparcos reference stars: they are captured during the very same time interval as the asteroid. The main benefit of this approach is that the reference stars are so close to the asteroid, typically a degree or so away, that it is possible to obtain good photometry under poor sky conditions especially where many images are taken over a longish time-period when relative differences in sky transparency tend to average out. Indeed, such an approach can also be used to perform absolute photometry on variable stars to high precision as described in the following section.

An alternative methodology is to use the ancillary scope to accurately measure the V magnitude of say an 11th or 12th magnitude star selected for its suitability as a comparison star in images taken with the main scope. Note that with this approach, there is no real need to use a check star since the actual V magnitude of the comparison is measured in real-time. However, it is always useful to use check stars since less images are required overall.

So far I have avoided the mention of the need to transform instrumental magnitudes to the standard system. Even with V-filter images, there will be a need to correct the instrumental magnitudes for differences in color of the asteroid and comparison star relative to those of the reference stars. Although the colors of the Hipparcos stars are known to a sufficient accuracy for this purpose, those of 11th or 12th magnitude comparison stars are probably not available. Similarly, where variable star photometry is contemplated, the variable can vary tremendously in color and the only way to obtain an estimate of its color is to measure this directly. If your main scope has a filter wheel then it is a relatively simple job to take images in V and B, or V and I at the start of an observing run, from which it is possible to select a comparison star similar in color to the asteroid (or variable). Currently, I do not have a filter wheel for

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Table 1. Absolute photometry. Measurement of V-band Extinction and Exo-atmosphere Zeropoints.
the main scope so instead I chose an alternative option: that is to add a second 60-mm refractor equipped with a CCD camera and I-filter. Images depicting the setup are shown below in Figure 4.

Using the two V- and I-scopes, I operate these simultaneously so as to improve the accuracy of the V-I color determination. I chose an I-filter for the second ancillary scope mainly because asteroids are significantly brighter in the I-band compared with the alternative choice of a B-filter.

I have written an Excel workbook to facilitate data reduction allowing for differential extinction and color transformation. As a test of the system, I carried out precision photometry of a dozen variable stars between 2004 September and 2005 March. The next section describes results obtained for a well-behaved red variable so as to illustrate the potential precision achievable for a stationary target with fixed reference stars.

5. Variable Star Photometry

Simultaneous V/I-photometry is a powerful tool for studying variable stars especially if one wants to follow changes in color of an object since fairly precise observations can be made during poor sky conditions. Similarly, if comparison stars are very close to the variable under study, then differential photometry can yield good results at times when sky transparency is fluctuating. For example, V370 Andromedæ was imaged on 30 nights, and measured against an ensemble of 5 Hipparcos stars. Several other field stars were also measured, such as HD 11904 which yielded a mean magnitude of V=8.146 for which the associated standard deviation was only 0.005 mag even under a variety of sky conditions.

A range of stars were monitored, some of which were Gamma Cassiopeiae-type variables, which showed a seemingly chaotic night-to-night variation of a few hundredths of a magnitude. Other stars were cool irregular red variables such as BZ Andromedæ, which were well-behaved and provided a good test of night-to-night photometric performance. A plot of the lightcurve of BZ Andromedæ based on 49 nights’ observations is shown in Figure 5, depicting four maxima and four minima over a six-month period. The entire variation in brightness is less than 0.5 mag yet virtually all of the individual datapoints lie on the line representing a smooth fit to the data. Clearly this result demonstrates again that high precision of the order of a few millimags is achievable night-to-night with the current hardware.

6. Clear-to-V Magnitude Photometry

The main 280-mm aperture telescope is utilized in unfiltered (Clear) mode so as to maximize signal to noise for faint asteroids. Incorporation of a V-filter would otherwise result in about a 75-80% reduction in signal. Operating unfiltered permits photometry on NEOs down to about V=18 with an absolute accuracy of about 0.05 mag.

As mentioned earlier, the fact that asteroids exhibit a relatively narrow range of colors compared to variable stars makes transformation from unfiltered ‘raw’ magnitudes to standard V magnitude entirely feasible. (N.B. This is not the case when it comes to measuring red or very red variable stars).

I shall explain one approach as to how Clear-to-V transformations can be carried out on asteroids with good precision. Again, I shall resort to using the Hipparcos dataset as a valuable resource permitting accurate calibration of unfiltered photometry.

So-called ‘red-blue’ pairs are often used to calibrate photometers: this is a well-accepted practice. Potentially, a large number of red-blue pairs can be gleaned from the Hipparcos dataset. My approach was to identify a total of 90 stellar pairs separated by <14 arcmin and distributed in a 5-degree wide band centered at Declination +20 deg. They have been selected according to a number of criteria. For instance they are relatively bright (6<V<10), not variable, not double (unless very close), do not have a nearby field star complicating the photometry, etc.
Figure 5. Slow irregular variable star, BZ Andromedae: Example of 49 nights of photometry.

Figure 6. Distribution of Hipparcos red-blue pairs.

The distribution of the selected pairs is shown in the diagram, Figure 6. Calibration is performed on a ‘photometric’ night by taking say 10 images of each red-blue pair as each approach culmination (airmass ≈ 1.2) using exposures times that avoid saturation of the CCD pixels. Typically a dozen pairs can be imaged in an hour or so. Each red-blue pair provides a measure of the transformation coefficient for the system. The conventional approach is to plot v-V, where v is the instrumental magnitude, against color for a set of stars in an open cluster such as M67, the slope of the plot giving the transformation coefficient, Tv. With red-blue pairs there are in effect just two stars to plot but each pair yield a measure of the slope, Tv. An estimate of the uncertainty in Tv is obtained by measuring several pairs and plotting Tv against color. If a V-filter is used, Tv should show negligible dependence on color, indeed it is assumed to have a zero second-order term in the conventional approach. However, an unfiltered CCD will show a very significant second-order dependence, a point that is not so widely known.

To measure the response of my 280-mm Schmidt-Cass telescope and unfiltered Starlight-Xpress SXV-H9 CCD camera, I imaged 13 red-blue Hipparcos pairs. The data reduction proceeds through several steps. The Hp magnitude is first converted to Johnson V using the following empirical equation based on the correlation between these parameters for stars measured by both Landolt and Hipparcos:

\[ V - Hp = 0.1238 \times (V - I)^2 - 0.2814 \times (V - I) + 0.0018 \]

This relationship is valid for the range, 0.0 < (V - I) < 2.0 (Correlation R² = 0.978).

The difference in V magnitude, ΔV of each pair of stars is then derived (in the sense blue minus red), typically accurate to better than 0.01 mag. The observed difference in the unfiltered instrumental magnitude, Δv is also calculated, from which the transformation coefficient is given by:

\[ Tv = (\Delta v - \Delta V) / \Delta (V - I) \]

Again, the precision of the Hipparcos stars is demonstrated by the small scatter in Figure 7: this shows a strongly color-dependent transformation coefficient as a function of the mean color of each red-blue pair. Using this unfiltered system, the scatter in individual measurements of Tv amounts to a mere 0.009, i.e. less than 1% standard deviation.
Figure 7. Plot of transformation coefficient versus color for an unfiltered system showing the strong second-order dependence on color.

The linear dependence of $T_v$ on color, as exhibited in the above plot, facilitates the transformation of unfiltered (Clear) magnitudes to $V$ magnitudes, since the correction term or degree of departure in instrumental magnitude from the standard $V$ magnitude is obtained by integrating the linear relationship shown in Figure 7 to yield a second-order polynomial expression as shown in Figure 8.

Here it can be seen that the correction is virtually zero for objects having a $V-I$ color of about 0.6-0.8. Of course the choice of CCD is an important factor if you wish to operate unfiltered yet preserve an accurate transformation to standard magnitude. The SXV-H9 camera equipped with a Sony ICX285AL Exview interline imager has a general response closer to the visual range than do many other CCD cameras, which tend to be more red-sensitive; so the latter type tend to be a closer match to the $R$ passband. In effect, when imaging objects having a $V-I$ color of 0.68, this unfiltered system has the same response as that of the standard $V$ photometric system. However, either side of this color (blue or red), the system tends to overestimate brightness relative to the true $V$ magnitude.

Applying the correction in practice, we need to know the color of the reference star and that of the asteroid (or variable). If the asteroid has a $V-I=+0.9$ say, and the comparison star has a $V-I=+0.2$, then the correction to the differential magnitude (in the sense, asteroid minus comparison) requires subtracting 0.022 mag (derived from the difference between 0.007 and 0.029 mag as obtained from the relationship plotted in Figure 8) from the measured value, $\Delta v$.

There is a further very important factor that needs to be taken into account when applying corrections to unfiltered magnitudes and that concerns the way in which the atmospheric extinction coefficient for the unfiltered passband has a significant dependence on the color term. This atmospheric factor is similar to the second-order extinction term that is applied when observing in the $B$ passband. Using a $V$ filter, the second-order extinction term is usually negligible but this is certainly not the case for unfiltered photometry. So the correction factor shown in Figure 8 is only strictly valid if observations are made at a similar airmass to that used in deriving the calibration.

Consider what the effect of moving to higher airmass would be. Here the atmosphere is more transparent to the red end of the spectrum than to the blue. So stars which are redder will appear to be brighter than they should be. I have carried out a series of observations of 10 red-blue Hipparcos pairs at different airmass so as to determine the second-order coefficient. The analysis was performed by calculating the difference in the frame zeropoint relative to the difference in color of the two stars in each pair. For a given color, stars at higher airmass tended to show greater differences in frame zeropoint values. Indeed, by dividing the measured difference in frame zeropoint by the mean airmass of each pair it was possible to arrive at a measure of the second-order atmospheric correction term as a function of the difference in color of the two stars as shown in Figure 9.
Here it can be seen that for this unfiltered system, the redder the star, the brighter it appears by a factor equal to 0.044 mag per airmass change per unit change in B-V. (Note that the values of B-V and V-I are in practice roughly the same, hence a similar correction would be necessary working in V-I). This second-order relationship was found to be valid up to an airmass of at least 10.

So to convert unfiltered magnitudes to V, we need to apply a correction based on the relationship shown in Figure 8 but which is adjusted to take into account airmass. Take for example Figure 10, which shows the resultant correction plot for airmass, X=3.0. In general, the full correction for transforming from Clear-to-V for this unfiltered system is obtained by substituting in the relevant values for the color index of the variable and comparison star in the following expression:

\[ y = 0.123*(V-I)^2 + k''*X*(V-I) - 0.221*(V-I) \]

where \( k'' \) is the second-order extinction coefficient, found to be 0.042 for the night when the data shown in Figure 9 were obtained. The difference in the value of the above relationship for the two objects, expressed in the sense variable minus comparison, is then calculated and subtracted from the measured differential magnitude to arrive at an accurate value of \( \Delta V \).

The second-order extinction term is very important when observing unfiltered and it should be realized that it will depend on the transparency of the night in question. In practice, \( k'' \) is likely to be dependent on the value of the first-order extinction coefficient, \( k' \). The first-order term is measured along the lines described in Section 4 of this paper using extinction stars selected from a set of 311 Hipparcos stars having B-V colors in the range 0.800-0.850. A convenient interactive Excel workbook has been designed which aids both the selection of stars in real-time, gives required exposure time for each, and has a calculation sheet which yields the extinction coefficient, errors, etc. when photometric adu values are pasted into it (using current AIP4WIN format).
7. Low Phase Angle Photometry

To further elucidate the light-scattering properties of asteroid regoliths, it is important to define the exact form of the opposition effect brightening beyond a simple linear change in reduced magnitude with phase angle. In some cases this can amount to a 0.4 mag increase in brightness, with the greatest increase occurring as zero phase angle is approached as depicted in the following Figure 11 taken from the article by Muinonen et al. (2002) published in the book, Asteroids III. The task of the observer wishing to carry out phase curve studies is to determine the exact form of these phase curves for as many of the various types and families of asteroid as possible.

During 2005 February to April, I began a campaign to determine phase curves for asteroids reaching very low phase angle, and managed to complete imaging runs on three asteroids over a wide range of phase angles, which should prove a good test of the observing methodology outlined in this paper. Full analyses of the data will be presented at the Society for Astronomical Sciences meeting at Big Bear, California on May 25th.

The first object observed was the asteroid 133 Cyrene, selected not only because it passed through and was observed at a phase angle as low as 0.11 deg, but also because this object was the subject of a careful study by Harris et al. (1984). In all, images were secured on 20 nights up to a phase angle of 17.8 deg. This object has a well-characterized rotational period of 12.708 hours (Harris 2005), knowledge of which facilitates the analysis.

A second object, asteroid 112 Iphigenia, has also been imaged on 18 nights at phase angles ranging from 0.22 deg to 18.6 deg. Preliminary results show it to have a rotational lightcurve amplitude of a little less than 0.20 mag and a period of more than 15 hours, making it a difficult object to characterize using conventional differential photometry. Harris (2005) indicates that this object has a rotational period of 15.783 hours and an amplitude of 0.5 mag. However, he ascribes the lowest reliability code of 1 to these values and notes that no lightcurve for 112 Iphigenia has been published. Clearly, this object is a useful challenge and will provide a good test of the observing methodology described here.

Finally, 17 nights of imaging were carried out with asteroid 130 Elektra as target, spanning phase angles of 0.33 deg – 14.9 deg. Harris (2005) indicates this object to have a rotational period of 5.225 hours and an amplitude of 0.19-0.58 mag depending on heliocentric longitude. He also notes that it is a suspected binary on the basis of a secondary, non-commensurate periodicity in the lightcurve, indicating possible mutual eclipse phenomena. Results from the present observing campaign should prove interesting.

8. Some Remarks and Conclusion

Characterizing asteroid phase curves is a much under-researched topic and the limited amount of good-quality data currently available inhibits theoretical modeling of the surface light-scattering properties.

Provided observers can migrate away from differential photometry and successfully master the challenge of absolute photometry, they can help to further elucidate the nature of asteroid regoliths via an improved knowledge of their light-scattering properties.

Automated surveys have made tremendous inroads in the field of asteroid observation, so much so
that discovery of new asteroids is almost beyond the reach of even the advanced amateur, and astrometry is only practical for NEO follow-up. Indeed, once PanSTARRS is operational, the sky will be mapped down to very faint limits (V=24) although many interesting objects will appear saturated in the survey images preventing accurate photometry. However, I do not expect the automated systems currently operating, or those in the pipeline, to seriously challenge the usefulness of amateur contributions to phase angle studies especially since amateurs working together can both obtain good observational coverage down to extremely low phase angles, and carry out extended imaging runs to define the shape and period of each rotational lightcurve.

During the coming months, it should be possible to establish whether or not quasi-absolute photometry as described here is a valuable observing methodology for deriving asteroid phase curves under mixed sky conditions. We shall see.

One topic not dealt with in the present paper concerns the problem of suitable software to facilitate the data reduction, which is the current rate-limiting step of the process. My view here is that a preferred observing methodology first needs to be established so that the question of how best to reduce the images, etc., can then be tackled. Automating the reduction pipeline would then be the holy grail!

9. Acknowledgements

I would like to thank E. Norman Walker for providing the high-quality V- and I-filters. The fact that they are sandwiched between 1-mm thick silica plates provides for good long-term stability.

Thanks also to Al Harris for pointing out his co-authored 1984 paper on 133 Cyrene, and for his marvelous encouragement over the past 20 years or so.

Like many others, I am also grateful to Brian Warner: in my case for establishing an area of his

Collaborative Asteroid Lightcurve Link (CALL) site which includes reference to a Low Phase Angle project, and in particular for providing a listing of objects approaching low phase angle, which is very helpful for identifying potential targets and planning observing schedules (Warner (2005)).

10. References


Archiving Lightcurve Data in the NASA Planetary Data System (PDS)

Mark V. Sykes,
David Tarico,
Rose Early
Planetary Science Institute
1700 E. Fort Lowell, Suite 106,
Tucson, Arizona 85719
sykes@psi.edu

Abstract

Ground based observers may submit data to the NASA Planetary Data System using a new web-based interface, the PDS On-Line Archiving Facility (OLAF). The PDS is a peer-reviewed data archive in which data are sufficiently characterized so as to be useful beyond the lifetime of the submitter. PDS data are given citations for reference in publications, and are widely distributed electronically. © 2005 Society for Astronomical Science.

1. Introduction

Ground based observations of solar system objects play a significant role in the development of mission definitions, science goals, and the interpretation of mission data. The NASA Planetary Data System has for years archived these data, particularly in the areas of asteroids, comets, rings and planetary atmospheres, but this has been limited by the difficulty users understandably have learning to create required support files conforming to PDS standards detailed in a several hundred page document, and the limited manpower available at PDS to make corrections to those files when submitted.

This problem has been overcome through the development of a web-based interface in which PDS standards are imbedded in its software so that the user need only fill out forms and upload a “tar” or “gzip” file containing their data. This is the PDS On-Line Archive Facility, or OLAF. OLAF is able to accept data in the form of ASCII tables or FITS images (no Word, PDF, or Excel, for instance). Specialized interfaces for common data types such as time-series or spectra are included.

2. Requirements

Time series data (e.g., lightcurves) are required to have three columns in an ASCII table - time, observable (e.g., magnitude), and uncertainty or error (optional). There should be only one table per target object. All the information about the observations is gathered in an index file (described in OLAF) or in the forms you are asked to fill out. Once submitted, the software checks for internal consistencies and, when accessed by the PDS, automatically generates all required PDS standards-compliant support files using the information you input. So less work for you, and less work for PDS.
3. Accessing OLAF

OLAF is accessible at

http://sbn.psi.edu/olaf/

Because archived data are expected to be of value for many decades or longer, all data and their ancillary information submitted to the PDS undergo peer review. This ensures that all information needed for long-term understanding and use of the data is included and that the data are usable. PDS datasets are now given citations (with information provided by the submitter).

Optionally, each data product (e.g., an individual spectrum or lightcurve) can also be given a citation. Citation information allows users and others to reference the archived data in the scientific literature. PDS is also transmitting citation information to the NASA Astrophysics Data System Abstract Service, which provides links directly back to the data in the PDS. Online and telephone support are available.

4. Acknowledgements

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Using a Distributed Observer Network to Characterize Transit Light Curves of Exoplanet TrES-1b

Ron Bissinger
Raccoon Run Observatory
1142 Mataro Court
Pleasanton, CA
ronbissinger@comcast.net

Abstract

Individual transit light curves of exoplanet TrES-1b were obtained and combined through a collaboration of amateur and professional astronomers functioning as a distributed observer network. Statistical and time series analytical tools were applied to a transit model, suggesting the presence of flux variations in and out of transit. Tools not commonly employed by amateur astronomers, such as bootstrap Monte Carlo and spectral signal analysis, were found to be readily available as off-the-shelf packages and useful for detecting and characterizing exoplanet transit light curves. The collaboration with professional astronomers on TrES-1b validates the important role amateur observers can play in conducting exoplanet research. © 2005 Society for Astronomical Science.

1. Introduction

Many amateur astronomers have measured the signature light curve dip of known transiting exoplanets such as HD209458b and TrES-1b. The increasing capabilities of their instruments and skills are allowing amateurs’ engagement, along with university and professional astronomers, in the search for new transiting exoplanets. As an increasing number of sub-meter aperture instruments become employed in these searches, it is important to quantify the detection levels that can be attained by such equipment and to identify analytical methods that can be used to detect and characterize transit light curves.

Wide-field sky searches to detect the signature light curve dip of transiting exoplanets rely on fairly complex, custom software employing detection algorithms that analyze hundreds of stars simultaneously. The targeted searches done by amateur astronomers, such as those undertaken by Transitsearch, often target a single star for a defined time window extending over several evenings. While transit depths of 1-2% (~10-20 mmag) similar to those of HD209458b and TrES-1b can readily be identified by visual inspection, current and future transit candidates will have more shallow transit depths. The capability to detect transits of 0.3% (3 mmag) or less, for example, will be required to identify systems with Jupiter-sized planets orbiting an A4 star, a Neptune-sized planet orbiting a K7 star, or an Earth-sized planet orbiting an M8 dwarf star.

An individual observation by itself, however, usually is not adequate to confirm a new transiting exoplanet, particularly if the transit depth is small. Instrumental noise, changes in observing conditions, and other sources of error conspire to inject doubt into any single observation. But if multiple observers pool their data and composite light curves are assembled, the noise in the data is theoretically reduced inversely proportional to the square root of the number of independent data sets. By coordinating multiple observers to generate light curves of a single known or suspected transit event, confidence can be significantly increased when the individual light curves are combined. Thus a coordinated observing network of amateur astronomers can generate light curves with higher precision than any single such observer, potentially approaching the precisions obtained by major ground-based observatories.

2. Transit Detection Capability of Amateur Telescopes

The observation of TrES-1b transits in 2004 provided an opportunity to gather light curve data from many amateur astronomers. Employing telescopes with apertures from 20 to 35 cm, experienced observers often achieved precisions of 3 – 4 mmag, particularly when multiple exposures are stacked over several minutes. Skilled amateurs also obtain similar precisions for other objects such as variable stars.

In order to quantify the transit detection thresholds of amateur-sized telescopes, a transit light curve model was constructed. Figure 1 shows the geometry of a transit event, including the 4 contact points and the pre-ingress, mid-transit, and post-egress segments.
A simple geometric model by Bruton (2005) with no limb darkening was used to model a generic transit event. Transits of exoplanets smaller than the hot giants like HD209458b and TrES-1b will likely exhibit not only smaller transit depths but longer transit times as well. If the transit times exceed a couple of hours it is probable that only the ingress or egress will be detected in a single observation. The transit model therefore used only the ingress with a depth that could be adjusted. The transit light curve model is shown in Figure 2.

The ubiquitous Microsoft Excel spreadsheet has spawned an entire industry for developers of add-in applications, and there are several sources of good, inexpensive Monte Carlo simulators, random number generators, and the like. It was fairly simple though to construct an Excel spreadsheet that did a bootstrap Monte Carlo with replacement for thousands of re-samplings, driven by macros. Simulated noise of various levels was added to the transit model using Excel’s random number generator. The randomness of Excel’s internal random number function was verified using a commercial Excel add-in statistical package, Crystal Ball. Crystal Ball used a separate, more robust random number generator to re-sample the Excel noise and confirm its random distribution (best fits were either Normal or Weibull distributions).

A metric in the bootstrap Monte Carlo is required as an indicator for the presence of a transit signal. The correlation coefficient $r^2$ between data sets proved not to be a reliable measure in the case of time series analysis as illustrated in Figure 4.

Assuming a periodic function like the simple sine wave shown in Figure 4, the correlation coefficient is strongly influenced by the slope of the data and not necessarily by any real differences between the data. Using $r^2$ in transit light curve analysis can be problematic if differential extinction is present or if there are irregular periods in the data.

A more accurate measure for comparing photometry time series data sets is the root mean squared (RMS) difference as it is less susceptible to influences from slope. Table 1 shows the results of the bootstrap Monte Carlo runs with the number of observers varied, showing the effects of stacking multiple independent data sets of the same transit. Each independent observer was assumed to gather data over 6 hours at 5 minute time resolutions (48 data points) with 4 mmag RMS noise, fairly conservative estimates of what should be attainable.
Transit Light Curves of TrES-1b – Bissinger

The results indicate that stacking light curves from at least 4 observers with typical noise levels and comparing them using a bootstrap Monte Carlo to the same number of stacked off-transit baselines can reveal the presence of transits with depths of 2 – 3 mmag. It is unlikely that a single observer, however, could identify a transit with a depth much less than 4 mmag.

Bootstrap Monte Carlo testing has limitations. It can’t characterize the transit signal by providing either timing or depth; it can only indicate the presence of a transit signal. The technique should not be applied to small sample sizes. A review of the literature suggests sample sizes of at least 20 to 30 should be used to obtain reliable results. The off-transit baseline data should also be obtained from the same observers with the identical instrumental configuration and under the same air masses and photometric conditions as used to obtain the transit data. Observations of different transit events using different equipment can introduce additional sources of noise and errors that will not be reduced by combining multiple light curves. Performing bootstrap Monte Carlo or other statistical techniques on such light curves can lead to uncertain results.

3. TrES-1b Discovery and Light Curve Database

In the summer of 2004 Alonso (2004) announced the discovery of exoplanet TrES-1b which was named after the network of small aperture telescopes (Transatlantic Exoplanet Search) used to detect the planet by the dip in the parent star’s light curve during transit. TrES-1b is located in the constellation Lyra at RA=19:04:09.8 (J2000); Dec = +36:37:57 (J2000), and orbits a K0V star with a magnitude of R=11.34, V=11.79, B-V=0.78 (Alonso 2004). A close-in giant planet, TrES-1b has a radius of 1.08 R\text{jup}, a mass of 0.729 M\text{jup} , an orbital period of 3.03 days and an orbital inclination of 88.2 degrees. Stellar parameters are R\text{star}/R\text{solar} = 0.83 and M\text{star}/M\text{solar} = 0.87 (Laughlin 2005).

Following the TrES-1b discovery announcement many amateur astronomers also recorded the exoplanet’s transit light curves and shared their results among such organizations as the American Association of Variable Star Observers (AAVSO) and Transitssearch (2005). Several observers (Gary, Garlitz 2004) noted that segments of the light curves leading up to ingress and leading away from egress seemed to exhibit brightenings.

![Figure 5. TrES-1b egress light curves from 120 and 35 cm telescopes; transits were approximately 3 months apart.](image-url)

Figure 5 shows one of the many apparent instances where similarities in flux variations from different observers and different transit events were apparent to visual inspection. Such strong visual correlations, however, need to be supported by more robust analysis and additional observations.

Therefore light curve data from amateur astronomers as well from the original discovery sources (Alonso 2004) were assembled. 34 separate observations of TrES-1b transits were available, totaling over...
4400 datapoints. The telescope apertures used ranged from 10 cm to 120 cm, and time resolution of the data was as short as 36 seconds and as long as 8.7 minutes. Including the original observations, the first transit was observed on May 9, 2004 and the last on October 11, 2004. Only 2 transits were observed at the same time by multiple observers, and then only for the mid-transit segment. Amateur astronomer CCD cameras used cover a wide swath, including SBIG ST-7XME, SBIG ST-8XME, SBIG ST-10XME, SBIG ST-1001E, and home made “cookbook” CCD cameras; their telescopes were on mounts from various manufacturers including Celestron, Meade and AstroPhysics.

For purposes of the current analysis, observations from instruments with apertures smaller than 20cm were excluded due to their reduced ability to resolve small flux variations both in intensity and in time. As shown in Table 2 the observations were made using a wide variety of time resolutions. It should be noted that the time resolution period for an exposure sequence consists of the time the CCD camera shutter is open, the download time of the image to the computer after the exposure is taken, a possible delay for the CCD camera to re-center a guide star, camera idle time and a multiplier if the images are stacked.

Table 2. TrES-1b Transit Observations

<table>
<thead>
<tr>
<th>Source</th>
<th>Telescope Aperture, cm</th>
<th>Location</th>
<th>Filter</th>
<th>time resolution, min</th>
<th>Transit Midpoint</th>
<th>number of datapoints</th>
<th>number of datapoints</th>
<th>number of datapoints</th>
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<td>V</td>
<td>1.70</td>
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<td>Canary Islands</td>
<td>I</td>
<td>1.68</td>
<td>174</td>
<td>24</td>
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<td>274</td>
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<td>1.68</td>
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<td>18</td>
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<tr>
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<td>51</td>
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</table>

Nyquist sampling indicates that time resolutions of 1 to 2.5 minutes or less are required to identify the short-period flux variations of approximately 3 minutes suspected to be in some of the light curves. Data having time resolutions > 3 minutes were excluded from the data used in this analysis as shown in Table 2.

4. Sources of Observational Errors

There were many potential sources of error in the observations as there was little consistency in the instrumentation and methodology used to obtain and process the images, perform photometry and generate the light curves. Airmass of the observations ranged from 1.0 to 1.7. Unfiltered observations were made as well as through V, R, r, g and z filters. At least one custom software application and three commercially available photometry software packages were utilized: AIP4Win, CCDSoft, and Mira AP. Each package may use different algorithms for centroiding stars and measuring flux during photometry. Different comparison stars were used in the photometry.

Significant errors are also likely in the reported timing of the observations. The time embedded in the raw CCD images is generally read from the internal clock of the computer controlling the CCD camera. It is not known whether all observers frequently synchronized their computer clocks using internet, radio or GPS technology; computer clocks frequently exhibit considerable drift during imaging sessions unless they are synchronized. Without frequent synchronization computer clocks can drift by tens of minutes over a night with certain software as image downloading interrupts the clock operation. If ob-
servers stacked images prior to running photometry there is also inconsistency in the time reported for the stacked image depending upon the photometry software used. Such inconsistencies can easily be on the order of several minutes.

5. TrES-1b Transit Model

In order to perform statistical analysis of the TrES-1b transit light curves, a model specific to the system was developed using the same geometric model developed by Bruton (2005) as was used for the transit signal detection exercise. The TrES-1b transit model included parameters from the TrES-1b system. Limb darkening was added using the equation \( a + b \cos(\theta + c) \), where \( a \), \( b \), and \( c \) are fitting constants and \( \theta \) is the scan angle across the diameter of the star. The base model with limb darkening was then visually fit to the observed light curve obtained by the 35cm CBA Belgium Observatory on JD2453250 using the previously published (Alonso 2004) 23 mmag transit depth. There is likely some dependency of transit depth on filtration due to limb darkening which was not modeled, and therefore a visual fit with a well-defined full transit light curve was judged adequate for the purposes of this exercise.

The transit lasts approximately 140 minutes from first to fourth contact. From the available data it was difficult to overlap the partial curves into a single, complete light curve because in most cases the transit depths differed or were difficult to determine. To simplify analysis the transit was therefore divided into the three segments defined below:

<table>
<thead>
<tr>
<th>Segment</th>
<th>Time Period</th>
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<tbody>
<tr>
<td>Pre-ingress</td>
<td>110 to 80 minutes before transit midpoint</td>
</tr>
<tr>
<td>Mid-transit</td>
<td>20 minutes before to 20 minutes after transit midpoint</td>
</tr>
<tr>
<td>Post-egress</td>
<td>80 to 110 minutes after transit midpoint</td>
</tr>
</tbody>
</table>

Table 3. TrES-1b Transit Segments

These time periods were chosen to exclude the steeply sloped shoulders of the ingress and egress, which would be difficult to analyze due to their large flux changes over short periods of time. The pre-ingress, mid-transit, and post-egress portions of the light curves were aligned both in flux and time by minimizing squares of differences.

Observed flux variations were then added to this base transit model using data from the multiple observations. The 35cm Racoon Run and 120cm F.L. Whipple observatory light curves presented the most distinct flux variations pre-ingress and post-egress and therefore were used to develop an empirical model for those segments. The model is symmetrical around both sides of the transit (pre-ingress and post-egress). Time shifts required to align the light curves ranged from 4.6 seconds to 5.5 minutes. The observed flux variations were modeled by fitting an equation consisting of eight sine terms as shown in Figure 6.

Flux variations were also visually apparent in the mid-transit segment in light curves made of the same transit by two different observers on JD2453280. The two observers, Racoon Run Observatory and Garlitz, used completely different equipment in different locations: California vs. Oregon; 35cm SCT vs. 20cm Newtonian telescope; commercial AP1200GTO vs. homemade belt-driven equatorial mount; and commercial SBIG ST-10XME vs. homemade Cookbook 245 CCD camera. The model for these mid-transit flux variations is shown in Figure 7.

The final TrES-1b transit model, including the visually apparent pre-ingress, mid-transit, and post-egress flux variations is shown in Figure 8.
6. Bootstrap Monte Carlo testing of the TrES-1b Transit Model

A bootstrap Monte Carlo can indicate the probability that the TrES-1b transit model signal is present in the observed data. Using a minimum requirement of 20 datapoints, according to Table 2, 18 out of 33 light curve segments met the criteria. The RMS difference between the transit model and the observations was used as a detection metric for the flux variations that were in the model, similar to the approach taken with the data in Table 2. Table 4 shows the results of the bootstrap Monte Carlo testing.

<table>
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<tr>
<th>Source</th>
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<th>Filter</th>
<th>time resolution, min</th>
<th>Transit Midpoint, JD2453000+</th>
<th>value of detection algorithm</th>
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<td>g</td>
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<td>2.70</td>
<td>186</td>
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<td>1.29</td>
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<td>286</td>
<td>101%</td>
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</table>

Table 4. Bootstrap Monte Carlo testing of TrES-1b Transit Model

Firm conclusions are difficult to draw from the data in Table 4 due to the limited number of observations that could be tested. Because of the uncertainties in the mean RMS differences computed by the bootstrap Monte Carlo, an RMS ratio of <115% probably reflects a negative detection; the higher the ratio, the higher the likelihood of a positive detection.

Note that the mid-transit segments for JD2453274 and 280 were the only segments observed by different observers at the same time. While the mid-transit segment model was developed using only data from JD2453280, relatively high detection metrics were found in 3 consecutive transits on JD2453274, 277 and 280. The pre-ingress and post-egress segments of the model were developed using data from the Racoon Run Observatory. Yet the pre-ingress transit model signal did not consistently yield high detection metrics when applied to the data from the same observatory from the 3 consecutive transits on JD2453283, 286 and 289. There may be some suggestion that the transit model signal is detected more consistently in consecutive mid-transit segments than the pre-ingress or post-egress segments.

7. Time Series and Signal Analysis of the TrES-1b Light Curves

Several methods useful for analyzing signals embedded in noisy data can be applied to exoplanet light curves such as those from TrES-1b. A commercial signal analysis package (AutoSignal™ from Systat Software, Point Richmond, CA), offering an extensive library of spectral analysis, noise reduction, and decomposition tools was used. Periodograms, essentially discrete Fourier transforms of the pre-ingress, mid-transit and post-egress data, were generated. The data was also examined for non-stationary periodic signals using continuous wavelet analysis.

AutoSignal™ generates periodograms from unevenly sampled data using the Lomb-Scargle algo-
Algorithm that was originally developed by astrophysicists. Periodograms can identify the presence of waves or harmonics in the light curves. Examination of periodograms of the TrES-1b light curves would show the presence of cyclical systemic noise from telescope mount tracking mechanisms, for example, or other signals with fixed frequencies. Periodograms generated by AutoSignal™ show critical limits (50%, 90%, 99% and 99.9%) that differ from the more common confidence level definition. A confidence level of 90% applies to a single data set wherein 10% of the data points would lie above that level due to random chance. A 90% critical limit as used by AutoSignal™ is the level of the largest peak in only 1 of 10 separate random noise signals due purely to random chance.

Figure 9 shows the periodogram for the pre-ingress observations listed in Table 1. Numerous small peaks are seen but only one at a frequency of 11.215 (period of 5.35 seconds) exceeds the 90% critical limit. A similar peak at a frequency of 11.203 is also seen in the Figure 10 periodogram for the Table 1 egresses. However, the peak in the egress periodogram does not attain the 50% critical limit, nor is it present in the mid-transit periodogram in Figure 11, suggesting that it is either random or due to short-term seeing effects.

![Figure 9. TrES-1b Pre-ingress segment periodogram](image)
Figure 10. TrES-1b Post-egress segment periodogram

Figure 11. TrES-1b Mid-transit segment periodogram
The TrES-1b light curve data was also examined for moving pulses or waves using the continuous wavelet and short-term Fourier transform tools available in AutoSignal™. No evidence of non-stationary signals was found in the data.

8. Mid-transit Segment Analysis

The limited bootstrap Monte Carlo results in Table 4 along with the visually strong apparent correlation between light curves taken of the same transit by different observers as shown in Figure 7 suggested additional examination was warranted of the mid-transit segment light curves.

With time resolutions longer than 4.5 minutes, data from the 61cm University of Colorado Sommers-Bausch Observatory (SBO) available from Alonso 2004 was not suitable for the small time period and flux variation analysis but could provide additional insight into the mid-transit segment. The SBO observed a TrES-1b transit on JD 2453180, producing the mid-transit light curve shown in Figure 12.

Mid-transit light curves were available for 2 periods: JD2453174, JD2453180, and JD2453186 using professional observatories with large aperture telescopes of 61, 80 and 120 cm, and JD2453274, JD2453277 and JD2453280 using amateur telescopes of 20 and 35 cm aperture. Despite the difference of over 3 months, there was some strong visual similarities in the light curves as shown in Figure 13.

It was also possible to note the time peaks occur in the flux near the transit midpoint for the 2 periods since observations were made of sequential or every other transit. The results are plotted in Figure 14.

For the 3 highest flux peaks captured by the large aperture instruments on JD2453174, 180 and 186, there appears to be a strong correlation of position with time. No such correlation was seen in the data for the transits on JD2453274, 277 and 280. If the moving flux peak was caused by a stellar surface feature such as a starspot, the slope of the line in Figure 14 can be used along with the known stellar properties to calculate how long it would take to rotate around the star, yielding a period of approximately 195 days. Whether 195 days represents a multiple of the star’s rotation rate or whether it is the drift rate of a stellar surface feature with the planet tidally locked with the star cannot be determined at this point.

9. Conclusions

A variety of analytical techniques have been applied to transit light curves of TrES-1b. Bootstrap Monte Carlo methods have shown that multiple amateur astronomers generating coordinated light curves of single exoplanet transits can detect transits with depths of 0.3% or even less, providing an opportunity to find Neptune-sized and smaller exo-
planets. Combined with other readily available signal analysis software, the bootstrap Monte Carlo method can allow the detection and characterization of exoplanet transit light curves.

While suggestions of structure in the pre-ingress and post-egress light curve segments were found, stronger indications for flux structure on the order of several mmag were seen in the mid-transit segments. Such structures could be explained by either features rotating in the stellar chromosphere or the occultation of starspots by the orbiting planet.

Because there was little consistency in the available data, future observations of single TrES-1b transit events by multiple observers will be required to confirm the presence of any flux variations.

10. Acknowledgments

The author wishes to thank Dr. Greg Laughlin of the University of California, Santa Cruz for his thoughtful comments and insights. Thanks are also due to Tonny Vanmunster, CBA Belgium Observatory, Landen, Belgium, and Joe Garlitz, Elgin, OR for the use of their data in this paper.

11. References


Abstract
The author will discuss the collaboration between professional and amateur astronomers in the study of Cataclysmic Variable stars. In addition two case studies, V442 Oph and FS Aur, demonstrating this collaboration will be given with emphasis on the photometric process, data format and data submission. © 2005 Society for Astronomical Science.

1. Introduction

The Center for Backyard Astrophysics (CBA), cba.phys.columbia.edu, was formed to harness many amateur astronomers and their telescopes to perform long time series photometry of Cataclysmic Variable (CV) stars. In order to continuously cover these stars members of the CBA have been solicited from around the globe at all longitudes and latitudes. This group has been augmented over the years by professional astronomers that have recognized the power of a network of telescopes with motivated owners that are willing to observe selected CV’s.

The end product of this network are light curves that, due to the multi-longitude nature of the observers, are continuous over many weeks, months and occasionally seasons. These long observing runs are easily able to tease out those periods that are near 24 hours that would otherwise be lost for single observers.

2. CBA History

The beginning of the CBA was the Laurel, MD basement of Dave Skillman in the 1970’s. He called himself the Center for Basement Astrophysics as play on words of a better-known institute in Cambridge, Massachusetts. He started with photomultiplier tubes and ultimately progressed to a CCD camera, which he made himself. He produced light curves of Cataclysmic Variable stars and determined some of their basic properties.

In 1980 Dave’s work came to the attention of Joe Patterson of Columbia University. The two of them collaborated successfully with Dave providing the long-term photometry and Joe augmenting Dave’s light curves with stints at larger telescopes and spectrographs. In those early years they published multiple papers on the Superhump phenomena that they observed in these stars.

In 1993 they were able to bring in Dave Harvey in Tucson, AZ, which now broadened their longitude coverage. Rapidly as more scientific papers were published that demonstrated the worth of this broad time coverage approach, more amateurs joined the group. Now a problem developed. Most of the members did not work out of their basements as Dave did. It was decided that a better name would be the Center for Backyard Astrophysics. Some of this early history appeared in Sky & Telescope: Skillman (1981), Skillman (1993), Patterson (1998).
3. The CBA Today

The CBA network today has over 48 contributing members with many more followers that contribute on selected objects. The world coverage ranges in longitude from 175° East to 155° West and in both hemispheres.

The typical member has a telescope of 10” to 14” diameter with a good drive to which is coupled a CCD camera. The majority of the telescopes are computer controlled to some extent with many having scripting capability that allows un-attended operation. The images that are collected throughout the night are analyzed with commercially available software and the results are posted to an archive maintained by one of the members.

4. The Process

4.1 Data Collection

The goal is time-series differential photometry of the target star. What is needed is a series of sequential images of the program star field using 1-2 minute integrations. The integration time should be chosen such that the variable, comparison and check stars are not saturated. Over a nights time this can amount to 100-200 images. This process is repeated over ~50 nights. Usually the images are collected un-filtered because with smaller apertures, every photon that can be collected is precious.

4.2 Data Reduction

Once these set of images are collected they must be reduced to differential photometry magnitudes. In an ideal world each observer would use the same Comparison and Check star. However, due to different fields of view from each telescope this is not always possible. Often during a long campaign, members will e-mail each other and discuss which stars are being used. In any event the Comparison star should be chosen that is nearby in the field and as blue as possible. Again watch that the Comparison and Check stars don’t saturate the CCD chip.

With the hundreds of images that will be collected over the night it is imperative that some form of software be used that can analyze these images in a consistent manner. Fortunately there are several packages out there that will do this for a reasonable price.

4.3 Data Format

The format of the submitted data is very simple: a list of time series differential (Var-Comp) photometry and Julian Date. A text file that looks like this:

3472.63286 –0.112
3472.63359 –0.121
3472.63433 –0.145

Sometimes people add a third column, which could be Check-Comp or air mass. Both are somewhat useful. Note that the Julian Date is truncated with no heliocentric correction. Time should refer to the middle of the integration, not the beginning or the end.

4.4 Data Submission

The data can be either imbedded in an e-mail or attached to an e-mail and sent to the New Data Archive: cba-data@cba.phys.columbia.edu. The subject line of the email should contain the name of the CV and perhaps the JD range of the observations or the date the observations were made.

5. Contacting the CBA

The CBA is always looking for new members. If you are interested in joining the network or even receiving the news e-mails, send an e-mail to info@cba.phys.columbia.edu. Tell us a little about yourself and your desired level of involvement.

6. Case Studies

6.1 V442 Ophiuchi

On June 12, 2002 those of us in the network received the following e-mail from Joe Patterson:

“… It’s high time to take up the cause on V442 Oph. Long runs, densely spaced at all terrestrial longitudes you can dream up (including you naughty North Americans out there) – that’s what we want to properly subdue V442”

Over the course of 42 nights, there were 203 hours of time series photometry provide by six astronomers.
Figure 2. The field of V442 Oph with the Variable, Comp and check star noted.

V442 Oph

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</tr>
</tbody>
</table>

Figure 3. Light curve of V442 Oph, lower curve, over one night. This represents 215 individual measurements taken at the Vermillion Cliffs Observatory (CBA-Utah).

Once the data from all observers was analyzed it was presented in a paper published in the Publications of the Astronomical Society of the Pacific, Patterson (2002). An exceptional aspect of the CBA publications is that observers that contribute data to the understanding of the CV are included as authors of the resulting publication.
6.2 FS Aurigae
The call for observations this season of this CV came on December 2, 2004. This dwarf nova, with an orbital period of 85.7 minutes, has been a CBA candidate previously. FS Aur’s strange behavior was the topic of a publication, Tovmassian (2003).

This recent request for observations was coordinated with both spectroscopic and Chandra observations. As yet the relationships between these have not been delineated. Perhaps another season of coordinated observations will clarify this CV.

Figure 4. The field of FS Aur with the Variable, Comp. And check star indicated.

Figure 5. The light curve of FS Aur, bottom curve, obtained at the Vermillion Cliffs Observatory (CBA-Utah)
7. Conclusion

The Center for Backyard Astrophysics has been in existence for over 30 years and has produced over 34 publications of its findings. The concept that small telescopes operated by dedicated amateurs can produce good science is proven again. The CBA network of astronomers have proven invaluable to the understanding of Cataclysmic Variable stars and they have proven to be dedicated and reliable in the quest for time series photometric data.

8. References


Solar Sail Orbit Determination from Ground Observations: 
A Proposed Professional - Amateur Collaboration

Mark S. Whorton
NASA Marshall Space Flight Center
EV42/Guidance, Navigation, and Mission Analysis
Huntsville, AL 35812
Mark.whorton@nasa.gov

Abstract
Solar sail propulsion systems enable a wide range of space missions that are not feasible with current propulsion technology. Hardware concepts and analytical methods have matured through ground development to the point that a flight validation mission is now realizable. Astronomical observations may play an important role in the flight validation of solar sail propulsion systems. Astrometric data and visual magnitude estimation has great potential for contributing to orbit determination, thrust performance verification, and optical model validation. This paper presents an overview of solar sail technology and proposes a collaboration between astronomical imagers and mission analysts for a flight validation mission. © 2005 Society for Astronomical Science.

1. Introduction

With very few exceptions, all spacecraft missions have been designed and conducted according to the principles established by Johannes Kepler in the early seventeenth century. Kepler had the genius to assimilate Tycho Brahe’s voluminous observational data into a radical new paradigm. No longer did the Greek notion prevail that the heavenly bodies moved in perfect circles. To make Tycho’s data fit, Kepler reasoned that planetary orbits were ellipses with the sun at a focus, which became the first of his three laws of planetary motion. Later, Isaac Newton brought mathematical formalism to Kepler’s description of planetary motion. Spacecraft missions today are still designed with Keplerian elements and Newton’s laws of motion.

Scientists often devise mission objectives that are difficult to accomplish with current state-of-the-art technology. Missions such as asteroid surveys, high inclination solar orbits, and comet rendezvous place enormous demands on a typical reaction-mass propulsion system. Other missions demand an entirely new class of non-Keplerian orbits. Exotic missions such as stationkeeping at artificial Lagrange points and orbits displaced from the ecliptic require a continual thrusting for the duration of the mission. These important missions cannot be achieved with conventional expendable propellants. Solar sail propulsion systems have the potential to meet these mission demands.

2. Fundamentals of Solar Sailing

Solar sail propulsion utilizes the constant pressure exerted by the sun’s radiation to push the sailcraft along its path. Solar photons transfer momentum to an object during a collision, much like billiard balls colliding on a pool table. A photon’s momentum is the product of its mass and velocity, and while the latter is quite large, the vanishingly small mass means the photon momentum is quite small. The combined effect of a large number of photons is required to generate an appreciable momentum transfer which implies a large sail area. And since acceleration is inversely proportional to mass for a given thrust force, the mass of the sailcraft must be kept to a minimum.

Figure 1 illustrates how the thrust force is utilized for propulsion. Incident rays of sunlight reflect off the solar sail at an angle $\theta$ with respect to the sail normal direction. For the model assuming perfect reflectivity, there are two components of force. The first is in the direction of the incident sunlight and the second is in a direction normal to incidence that represents the perfectly reflected photons. When the two components vectors are summed the result is that the two components of force along the sail surface cancel each other and the components normal to the surface add together to produce the thrust force in the direction perpendicular to the sail surface. For a 40 meter x 40 meter square sail at 1 AU from the sun, the solar radiation thrust force is 0.0296 Newtons.
Solar radiation pressure can be used to either increase or decrease the orbit energy. If the sail is oriented such that the thrust force is opposite the direction of motion, as in Figure 1 for a heliocentric orbit, the orbit spirals inward. Conversely, if the thrust is in the direction of motion, the sailcraft orbit spirals outward. Orbit inclination changes result when a component of the thrust force is oriented perpendicular to the orbit plane.

Various configurations have been proposed for solar sail vehicles. One of the earliest concepts was a Halley’s Comet rendezvous mission using a heliogyro (middle of Figure 2). Heliogyros have reflecting surfaces formed by long blades rotating about a central axis and pitch controlled like a helicopter to provide attitude control. Another approach that is currently in development is a square sail comprised of four triangular sail quadrants. These systems are typically three-axis stabilized, as opposed to spinning, with attitude control torques provided either by articulating tip vanes or varying the center of pressure/center of mass offset. Spinning disk sails have also been proposed.

**Figure 1. Solar Radiation Thrust Force (NASA/JPL).**

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**Figure 2. Solar Sail Design Concepts (NASA/JPL).**

### 3. Solar Sail Flight Validation

For the past several years NASA has been investing in several aspects of solar sail propulsion technology. Teams have conducted research on analytical methods for modeling the shape and motion of the sail and support structures under solar radiation pressure loading. Others have developed prototype solar sail hardware systems including the sail membranes, structural supports, deployment mechanisms, and control actuation systems. This work has been converging toward a ground test of the prototype 20 meter x 20 meter square solar sail systems and the correlation of test data with analytical tools. Verifying the design processes and analytical tools on a scale system is a step in the process of validating the technology in a larger scale flight experiment. The first generation of science missions call for 80 meter square sails, so the intermediate goal for a flight validation mission in earth orbit is likely to be a 40 meter square sail.

A fundamental question is how one would go about verifying the performance of a solar sail as a propulsion system. The resultant thrust produced by a solar sail is a function of many variables that are difficult to measure on the ground or predict from analysis. The parameters can be loosely grouped into two categories: one associated with the thrust vector direction (i.e. pointing orientation of the sail) and the other associated with the thrust vector magnitude (i.e. reflectivity of the sail). Factors such as sail shape, both global and local; support structure deformations; disturbance torques such as atmospheric drag and gravity gradient; and the spacecraft attitude control system all influence the direction of the thrust vector. Likewise many factors such as membrane topology (wrinkles, billow, etc); specular and diffuse reflectivity; absorptivity; emissivity; and sail temperature affect the magnitude of the resultant thrust.

Two approaches may be used to estimate thrust performance – indirect methods and direct methods. Indirect methods utilize ground tracking station data and Global Positioning System (GPS) measurements for orbit determination from which the thrust profile that results in the measured trajectory is estimated. Orbit determination is the process whereby the spacecraft position and velocity is obtained, and in some cases also estimates errors in the analytical model of the spacecraft and measurement systems. Another approach would be to utilize acceleration measurements to directly measure the thrust performance. Enabling both approaches requires a navigation system architecture comprised of direct inertial measurements aided by GPS and ground station tracking data.

Several issues complicate the orbit determination and thrust estimation problem. Environmental forces, torques, and instrument biases all contribute to errors in direct thrust measurement. GPS measurement accuracy is a function of the orbit (altitude primarily) and hence may be useful for only certain times. Ground tracking data is a function of the availability, cost, and location of the tracking station. Hence a range of measurement data types combined with a judicious choice of orbit and attitude profiles is nec-
necessary to potentially isolate and eliminate error sources; model the sailcraft properties that affect the thrust vector; and accurately estimate the thrust performance.

4. Pro-Am Collaboration

One additional data source that can potentially be of significantly utility for thrust estimation is ground astronomical observations. Two types of ground observations are particularly beneficial: astrometric data and visual magnitude data.

Astrometric data is typically used to estimate the six constant Keplerian elements of satellite, comet, and asteroid orbits. In the case of solar sails, the Keplerian elements are functions of time. Hence the aggregation of astrometric data will be beneficial to reconstruct the trajectory and estimate the thrust vector time history. Likewise visual magnitude data will be useful to correlate spacecraft optical properties as a function of vehicle attitude. It bears repeating that these properties directly affect the sail’s performance as a propulsion system and are quite difficult to accurately predict or test on the ground. This astronomical data has the potential to directly impact the quality of the thrust estimation and model validation and hence may significantly contribute to the success of a solar sail flight validation mission.

The orbit chosen for flight validation is a compromise between affordable launch access and acceptable environmental constraints. One orbit under consideration is a 1500 km circular near-polar orbit. This orbit is high enough to effectively eliminate atmospheric drag, minimize gravity gradient disturbance torques, and guarantee no eclipses. Another alternative is a highly elliptical orbit with longer dwell times at apogee during which the effect of solar thrust can be maximized.

Astronomical imaging for mission analysis in either the circular or elliptical orbits will be a challenging endeavor. For the circular orbit, the angular velocity in the orbit plane is 2.98 degrees/sec and the angular size of the sail is 55 arcseconds. Figures 3 and 4 show the angular rate and size of the sail, respectively, as a function of apogee height in representative elliptical orbits. This is a much smaller angular rate than the circular orbit, but nonetheless this rate may pose a challenge for accurate tracking.

More than the tracking rate, the greater challenge will be contrast ratio between the bright solar sail and background stars. Several factors influence the visual magnitude: the angle of the sail with respect to the sun (“cone angle”); the angle of the sail with respect to the observer’s line of sight (“look angle”); the altitude of the sail; atmospheric seeing; and the sail’s optical properties among others. For a 1500 km circular orbit, Figure 5 shows the visual magnitude of a sail with 60% reflectivity for various look angles and cone angles. Figure 6 shows the visual magnitude as a function of altitude for the elliptic orbits with the sail oriented perpendicular to the incident sunlight. Obviously this corresponds to the sail at opposition and all look angles are not feasible, but this represents a limiting case.
partners who will work together to define the processes and infrastructure required for this task. This concept definition will then form a component of a flight validation mission proposal, which if approved, will lead to this concept being implemented as part of a solar sail flight validation mission.

Any comments or suggestions are welcome and may be addressed directly to the author.

6. References


5. Where We Go From Here

NASA, academia, and industry are working together to advance the technology of solar sail propulsion systems with the goal of validating the technology in flight. An important part of that flight validation will be the gathering of data to support the analysis for thrust estimation and model validation. A collaboration between the astronomical imaging community and the mission analysis team will be a synergistic effort that may potentially advance the state of the art for both communities.

Several key challenges remain to be addressed if this collaboration is to yield useful data. Issues such as equipment requirements, observation requirements and methods, observation times, image integration times, and the collaboration infrastructure remain open. We solicit input and hope to dialogue with the astronomical imaging community to address these issues and form a mission analysis team. Our goal is to establish a collaborative effort with interested
An Experiment in Relating CCD Differential Photometric Precision to Varying Degrees of Image Focus

Erick Sturm
Physics Department
California Polytechnic State University
1 Grand Ave, San Luis Obispo, California 93407
esturm@atl.calpoly.edu

Abstract
Obtaining precise differential photometry of variable stars, quantified by the standard deviations of the comparison-check (C-K) differentials, is vital in several types of variable star research. Many smaller observatories utilize comparatively low-cost, off-the-shelf telescopes and CCD cameras for their differential photometry. Relatively little research has been reported on the effect of focus on photometric precision for such observatories. A Meade 10" LX200 telescope and SBIG ST-7XE camera were used to take 15 one-minute images of a star field at five different degrees of focus. A pair of similar magnitude stars was used to form the C-K differentials for each of the five image sets. This process was repeated for other fainter pairs. Ultimately, this experiment may determine how the differential photometric precision improves or deteriorates for varying degrees of focus, and if that change is dependent on the magnitude of the star-pairs. © 2005 Society for Astronomical Science.

1. Introduction
Obtaining high precision differential photometry is vital in several types of variable star research. Normally the required photometric precision is achieved through gathering high numbers of photons in large-diameter telescopes equipped with custom, liquid-nitrogen-cooled CCD cameras, which decrease CCD thermal noise. Yet, many smaller observatories are financially limited to using smaller telescopes equipped with off-the-shelf, thermoelectrically-cooled CCD cameras to conduct their photometric research.

Relatively little research has been reported on the effect of focus on photometric precision for smaller scale observatories. Larger scale projects, such as the French COROT mission and the Kepler space mission, are proposing to purposefully defocus their images, although there has been very little documentation on this tactic (Howell 2000).

With respect to photometric precision, defocused images are expected to decrease the effect of intrapixel and inter-pixel variations, lessen the need to deal with partial pixels, and reduce the number of photons hitting a given pixel, thus allowing for longer exposure times and increased collection of source photons before reaching saturation (Howell 2000). Perhaps most important to smaller observatories, defocused images may provide a means of improving photometric precision without any additional monetary costs.

This particular experiment focused (no pun intended) on improving differential photometric precision (DPP) through varying the degree of image focus. The DPP was numerically quantified by the standard deviation of the comparison-check (C-K) differentials, denoted as $\sigma_{C,K}$. $\sigma_{C,K}$ is related to the standard deviation of the program-check differentials $\sigma_{V,C}$, as caused from noise and not the program star’s true variability, by a scaling factor $\Gamma$ as shown below (Howell 1988):

$$\sigma_{C,K}^2 \times \Gamma^2 = \sigma_{V,C}^2$$

Therefore, minimizing $\sigma_{C,K}$ leads to a corresponding minimization of $\sigma_{V,C}$, and thus a maximization of the DPP.

The motivation for this experiment comes from my involvement with the Orion and Dark Ridge Observatories and their differential photometric study of W UMa eclipsing binaries. Genet and Smith (2005) are working on determining the timescales of the changing photometric light-curve asymmetries and season-to-season rotational period changes of short-period W UMa-type binaries. Both of these require the utmost in photometric precision; thus the ultimate goal of this experiment to improve DPP.

In Sections 2 and 3 the data collection and reduction procedures are described. Section 4 presents the results of the experiment. Then, in Section 5 some additional considerations concerning the results of the experiment are discussed.
2. Data Collection

Data was collected on two occasions. The first run, on January 20, 2005, used a Meade 10” LX200 telescope equipped with a SBIG ST-7XE CCD camera. A Meade 14” LX200GPS telescope equipped with a different camera of the same model was used for the second run on April 9, 2005. All the images collected during both runs were captured through an R$_c$ filter.

During the first run, 15 one-minute images were taken of a star field centered on V523 Cas at five different degrees of focus. At the start, the micro-focuser was set such that it minimized the full width at half maximum (FWHM) of the point spread function (PSF) of the stars in the image. Adjustment to the next four focus settings was achieved by running the micro-focuser for five seconds from the previous setting. These focus positions were later quantified by the determining FWHM they produced in their corresponding 15 images. A sample image from each focus setting can be seen in Figure 2.

There was a fifth focus setting during the first run; however, at that point the image was so defocused that the autoguiding CCD lost its guide star, Figure 1 shows the resulting image. So, the fifth focus setting was thrown out and not reduced.

There was a fifth focus setting during the first run; however, at that point the image was so defocused that the autoguiding CCD lost its guide star, Figure 1 shows the resulting image. So, the fifth focus setting was thrown out and not reduced.

![Figure 1. Star streaks resulting from the autoguider's loss of its guide star due to extreme defocus.](image)

![Figure 2. Sample images from different focus setting where (a) has the lowest FWHM at just above 3 pixels and (d) has the highest at around 20 pixels.](image)
The second run used a star field centered on M67. Since a larger diameter telescope was being used, the integration time was decreased from 60 to 30 seconds. Also, the number of images at each focus position was increased from 15 to 30 exposures. Additionally, the experiment was conducted somewhat differently. The micro-focuser was set to produce a FWHM of about 10 pixels at the start. The subsequent focus positions were set such that the FWHM was reduced by approximately one pixel from the previous setting. This process was repeated until the best possible focus was reached, giving a minimum FWHM.

3. Data Reduction

The images were reduced using Software Bisque's CCD Soft, and the astrometric solutions were found with the combination of CCD Soft and The Sky 6. The photometric data was extracted by surrounding the center of the star with a circular annulus of a specified diameter. The center of the star was found assuming a Gaussian PSF. At the time this paper was submitted, only the images from the first run had been reduced, so all data extraction and analysis was done on only the data from the first run.

For the first run, the photometric data was extracted several times from the same set of images by varying the pixel diameter of the annulus in order to find an optimum annulus for two different magnitude star-pairs at the different FWHM values. These pairs of similar magnitude stars were then used to form the C-K photometric differentials. The star-pairs used in the experiment are shown in Figure 3. Table 1 shows the properties of both stars in each pair. The FWHM in Figure 3 is the smallest of the observing run at just above three pixels.

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</table>

Table 1. Properties of the stars in the pairs chosen to form the C-K differentials.

Once the data was reduced, $\sigma_{\text{C-K}}$ of each magnitude pair was then plotted against the diameter of the extraction annulus for each degree of focus (see Figure 4).

The extraction annulus coinciding with the lowest point on each curve in Figure 4 was considered the optimum annulus for that given combination of star-pair magnitude and FWHM. The optimum extraction annulus agreed well with the results of Howell (1989), where he finds the optimum annulus to be near one FWHM. Figure 5 shows the same data as in Figure 4 but the abscissa now reports the extraction annulus as a multiple of FWHM.
Photometric Precision & Image Focus - Sturm

As can be seen from Figure 5, the optimum extraction annulus diameter is about one FWHM for all the different FWHMs for magnitude pairs except the lowest, most sharply focused, one. The lowest FWHM has an optimum annulus of about three FWHMs in diameter. One possible explanation is that the PSF of the lowest FWHM is very close to the true Gaussian PSF that the extraction software is using to approximate the actual PSFs, making it a near perfect PSF according to the software. Howell (2000) suggests that the light from a star with a perfect PSF can be almost completely captured by an extraction annulus equal to three times the FWHM.

It should also be mentioned that even at the highest degree of focus (smallest FWHM) the data is still considered well-sampled. A sampling parameter \( r \) can be defined as:

\[
r = \frac{\text{FWHM}}{p}
\]

where \( p \) is pixel size, and FWHM and \( p \) are in the same units. For data to be considered well-sampled, \( r \) must be greater than 1.5. For \( r \) values around and less than 1.5, the data is considered marginally sampled and under-sampled, respectively (Howell 2005).

For this experiment, since the FWHM is reported in pixels, \( r \) is equal to the FWHM. Therefore the lowest \( r \) value is larger than 3.0 and as such the data is well-sampled. Thus, no special techniques of dealing with under-sampled data (see Howell 1996) were used in the reduction process.

After the data was reduced, the lowest point on each curve in Figure 5 was then used in the final comparison between \( \sigma_{C-K} \) and FWHM.

4. Results

Figure 6 shows the results of the experiment. The \( \sigma_{C-K} \) of the 13th magnitude star-pair is an order of magnitude higher than that of the 11th magnitude pair. According to Howell (1995), any “good” software used for data reduction can be expected to provide “good” photometric data as long as brighter sources are used. Therefore, the higher-magnitude star-pairs were expected to show a less clear relationship between focus and DPP since they were captured in the same image as the brighter, lower-magnitude stars and, consequently, are dimmer. The same is also true at the other end of the spectrum, where the brightest stars in the frame could fill the pixel wells to the point of non-linear response.

![Figure 5. C-K standard deviation plotted against the extraction annulus used as a multiple of FWHM.](image)

![Figure 6. The effects of varying degrees of focus on C-K standard deviation.](image)
lish a true minimum. The second observing run tried to meet this need; once the data is reduced a clearer trend between FWHM and DPP may be seen.

5. Considerations

The results obtained above vary based on the chosen stars in a field. Though a certain FWHM may be shown above to minimize $\sigma_{C-K}$, conditions may exist where this FWHM doesn’t produce superior results. If the star field is crowded, such that the chosen stars in the pair have very near neighbors, the PSFs of neighboring stars may overlap, leading to an increase in $\sigma_{C-K}$.

Another consideration is that of magnitude and signal-to-noise ratio, S/N. Although the 13th magnitude star-pair produced data that appeared to be dominated by noise, this is not necessarily a result of the magnitude of the pair but could instead be the S/N of the data. Since the images integration times were set for around 10th magnitude stars, the 13th stars had very low S/N ratios. If the experiment were repeated with longer integration time, the 13th magnitude star-pair might be expected to produce a result more akin to that of the 11th magnitude pair with the current integration time.

Another consideration is photometric under-sampling. If a system is used where the lowest FWHM yields an r less than 1.5, the under-sampling techniques discussed by Howell (1996) need to be employed in order verify that the results are indeed due to degree of focus and not poor photometric sampling.

6. Conclusion

The results suggest that the experimental procedure has the possibility of providing a relationship between $\sigma_{C-K}$ and FWHM. However, continued work is necessary if a solid relationship between $\sigma_{C-K}$ and FWHM is to be established. Since the second run used larger sample sizes at each focus setting, along with decreased step size between focus settings, it should establish a meaningful trend and provide an optimum FWHM for a given star in a given star fields once the data is reduced and analyzed. Reporting the different star-pairs in S/N as opposed to magnitude may prove to be helpful in making the results of the experiment more easily applied to different systems with various integration times.

7. Acknowledgements

I would like to thank Russell Genet and Thomas Smith for allowing me telescope time at the Orion and Dark Ridge Observatories, and for all of their encouragement and support.

Also, many thanks to John Mottmann for being my official research advisor at California Polytechnic State University. I would also like to recognize the Physics Department of the California Polytechnic State University at San Luis Obispo for their financial support of my senior research project.

And of course, my thanks to Brian Warner and the Society for Astronomical Science for editing and publishing my first paper!

8. References


Uncertainty Analysis in Photometric Observations

Michael Koppelman
The University of Minnesota

Abstract
Uncertainty analysis is an important part of any scientific measurement. We take a brief look at formal uncertainty analysis for any observation set, including systematic and random sources of uncertainty. We then address sources of uncertainty in photometric observations and discuss practical methods for applying uncertainty analysis to common photometric data sets.

1. Introduction

As any introductory physics class will emphasize, any measurement, to be useful in science, must be accompanied by an estimate of the uncertainty in the measurement. Astronomy often deals with measurements with extremely large numbers and as such can be forgiving when it comes to uncertainty. It is not uncommon to see estimates that are plus or minus an order of magnitude. Perhaps this gives us some insight into why astronomers are historically not very good at putting error bars on their plots.

Specifically on the topic of light curves obtained photometrically, one rarely sees error bars. There are several specific reasons why light curves so often go unadorned with error bars:

1. If the uncertainty is very small, often the size of the symbol used to plot the data point is bigger than the error bars;
2. If a comparison star or differential K-C is plotted on the same plot as the light curve, one can “eyeball” the uncertainty based on this information;
3. The author of the light curve did not calculate or did not plot the uncertainty.

Reason #1 above is forgivable, although this information should be clearly indicated on the plot. This is easy to do and provides the reader with all the information they need to have some confidence in the measurements. Reason #2 above is better than nothing but implicitly communicates the uncertainty rather than doing so explicitly. It is reason #3 above that is not acceptable. The uncertainty should be plotted, made available, or summarized. Otherwise, the data set is incomplete and has diminished value.

The purpose of this paper is to provide some mathematical background to the subject of uncertainty analysis and propagation and to provide some practical approaches for calculating the uncertainty of individual as well as time-series photometric measurements.

2. An Illustration of the Problem

Figure 1 shows a plot of a fabricated light curve. There is an apparent eclipse that we are interested in. Unfortunately, there are two serious problems with this plot. First of all, there are no error bars. Second, the plot has a line connecting the data points, implying that the behavior between the points is known.

![Figure 1](image1.png)

Figure 2 shows the same data without the connecting lines. Still no error bars, though. What should we make of that eclipse?

![Figure 2](image2.png)
Finally, Figure 3 shows the data with error bars. Our excitement about a possible eclipse has just vanished. Instead, we conclude, clouds, bad tracking or some other mishap has most likely contaminated a couple of data points. The fact of the matter is, we were shown the data but not the uncertainty and we were misled as a result. Uncertainty determination and error bars are a vital necessity when communicating scientific data.

3. Signal-to-Noise Ratio and Uncertainty

There are two common ways to calculate uncertainty in photometric measurements: with the signal-to-noise ratio (SNR) or with the standard deviation of measurements of non-varying stars in the field.

If the noise is dominated by the Poisson noise from the star being measured, the uncertainty is given simply by:

$$\delta q = \sqrt{N}$$  \hspace{1cm} (1)

...where N is the number of electrons (or photons) in the measurement. It is important to note that this does not work for analog-to-digital units (ADUs) – your counts must be in electrons. You can calculate the number of electrons by taking the ADUs times the gain of your CCD camera.

It is fairly simple to see that this also gives an estimate of the SNR ratio:

$$\frac{S}{N} = \frac{N_*}{\sqrt{N_*}} = \sqrt{N_*}$$ \hspace{1cm} (2)

For example, if the net counts, measured in electrons, for a given star was 64,000 the uncertainty is +/- 253 counts and the SNR is 253. In this case the uncertainty is usually stated as 1/SNR or 0.004 in this example. This is a *fractional uncertainty* in that it is the noise divided by the signal. Thus, the uncertainty in this example is 0.4%.

It would seem at first blush that this should solve all of our problems with uncertainty in photometric measurements, since what we do, in essence, is simply count the number of electrons in our signal from our CCD chip. In practice, it is not nearly that simple. Equation 1 assumes that the only source of random error is the counting statistical error itself. In reality there are many other sources of error including scintillation noise, tracking and guiding errors, changes in seeing or transparency, thin clouds, noise from read out, etc. The fact of the matter is, even with a non-variable star on a clear and steady night, the number of counts tend to vary beyond that indicated by Equation 1, pointing to other sources of errors.

Steve B. Howell in “The Handbook of CCD Astronomy” (Howell 2000) gives a more comprehensive formula for the SNR. This equation takes into account the Poisson noise as above but also takes into account the uncertainty from the background estimation ($n_B$), the noise from the sky ($N_S$), the dark current ($N_D$), the read noise ($N_R$), the gain ($G$) and the 1-sigma introduced from the A/D converter. Here $n_{pix}$ is the number of pixels in the aperture in question.

$$\frac{S}{N} = \frac{N_*}{\sqrt{N_* + n_{pix}(1 + \frac{n_{pix}}{n_B})(N_S + N_D + N_R^2 + G^2 \sigma_f^2)}}$$ \hspace{1cm} (3)

In practice Equation 3 will decrease the SNR and increase the uncertainty by several percent.

For a single measurement, Equation 3 does a pretty good job of estimating the uncertainty. For time series photometry, though, there are new sources of random errors to account for. A prime example is the automatic centroiding of the aperture. Differences in tracking and seeing from frame to frame can cause differing amounts of light to be captured by the ap-
ertures. Equation 3 above cannot take these sources of uncertainty into account. A much more simple practice is to use several measurements of constant stars to estimate the uncertainty in the measurements of the program star. A comparison star, usually designated C and a check star, K, provide a simple measure of uncertainty by taking the standard deviation of a series of K-C measurements. The standard deviation, shown here in Equation 4, calculates a sort of average deviation from the mean.

$$\sigma_{K-C} = \sqrt{\frac{1}{N-1} \sum (x-\bar{x})^2}$$  \hspace{1cm} (4)

A minimum of three measurement are needed to meaningfully determine the standard deviation and in practice this method works best for time-series observations containing many measurements. The beauty of using Equation 4 is that, in theory, it encapsulates the entire uncertainty of the measurements. As such, this is a very common method for determining uncertainty in photometry for time series observations. It does, however, leave room for improvement, specifically because it assumes that the program star is of roughly the same magnitude as the comparison and check star. We will discuss a slight improvement in Section 5 that addresses this assumption but first we address the propagation of errors.

### 4. Error Propagation

The most general equation for propagating uncertainty in a function of several variables is:

$$\delta q = \sqrt{\left(\frac{\partial q}{\partial x_1}\delta x_1\right)^2 + \cdots + \left(\frac{\partial q}{\partial x_N}\delta x_N\right)^2}$$  \hspace{1cm} (5)

Equation 5 (Taylor 1982) allows you to figure out the uncertainty in a function $q$ when $q$ is a function of more than one variable. This is very useful if you are doing any transformations of data. For example, suppose you have determined your instrumental values for $b$ and $v$ and their respective uncertainty, $\delta b$ and $\delta v$. You now want to calculate $b-v$ and its uncertainty. In this example our function $q$ looks like:

$$q = b - v$$  \hspace{1cm} (6)

The partial of $q$ with respect to $b$ is 1 and the partial of $q$ with respect to $v$ is 1 so Equation 5 above now looks just like we would expect:

$$\delta q = \sqrt{(\delta b)^2 + (\delta v)^2}$$  \hspace{1cm} (7)

The errors add in quadrature. Because the quantities are squared, it doesn’t matter if you are adding or subtracting values, the uncertainty is simply the square root of the sum of the squares of the individual uncertainties.

Another example is converting magnitude units to flux units. According to Zombeck (1990) one can calculate the log of the flux in the V-band with the following:

$$\log f_\lambda (m_v) = -0.4 m_v + \log f_\lambda (0)$$  \hspace{1cm} (8)

...where the last term on the right, in this case, is $3.64 \times 10^{-9}$ erg cm$^{-2}$ s$^{-1}$ A$^{-1}$. If you are converting a V-band measurement to a flux you also need to compute the uncertainty. For example, a star with magnitude $V=12$ corresponds $\log(F_V)$ of $-13.239$ erg cm$^{-2}$ s$^{-1}$ A$^{-1}$. If our uncertainty in the V-band was 0.02, what is our uncertainty in these new units? Equation 5 to the rescue. First, we have to calculate the uncertainty due to Equation 8, which is simply $0.02 \times 0.4 = 0.008$. Now our equation for $q$ becomes:

$$q = 10^x$$  \hspace{1cm} (9)

...where $x$ is the result from Equation 8 and is a log. To compute the flux from the log of the flux we simply take 10 to the appropriate power. Our equation for the uncertainty becomes:

$$\delta q = (10^x \ln 10) \delta x = q \ln 10 \delta x$$  \hspace{1cm} (10)

Plugging in the numbers we get that $V=(12.00 +/- 0.02)m$ corresponds to $F_V=(5.77 +/- 0.1) \times 10^{-14}$ erg cm$^{-2}$ s$^{-1}$ A$^{-1}$.

In many cases, the calculation of uncertainty is a several step process that involves propagating uncertainties through all transformations from the raw data through to the final reduced result.

### 5. A Practical Recommendation

While this math stuff is all very interesting and indeed useful for many scientific practices, the point of this paper is to address a practical approach for the common case of variable star photometry. This recommendation is summarized in Equation 11:
\[
\delta q = \sqrt{\sigma_{K-C}^2 + \left(\frac{1}{SNR}\right)^2}
\] (11)

Equation 11 adds in quadrature the uncertainty for the overall time series of the K-C measurements and the uncertainty from the SNR calculation. The standard deviation of K-C provides a single uncertainty estimate for every data point. The 1/\(SNR\) provides a unique uncertainty estimate for each point, but excludes many sources of noise. Together this approach has two advantages over using a single one of these uncertainty estimates alone. First and foremost, it clearly illustrates data points with low SNR. This can occur when thin clouds roll through or during intervals of bad tracking. Second, it accounts for differences in brightness between the program stars and the C and K stars. If C=K=12m, for example and the program star is at magnitude 14, the standard deviation of K-C vastly underestimates the error.

In cases where C=K=V, Equation 11 does overestimate the uncertainty a bit. In practice, though, one term or the other usually dominates and the overestimation is slight. In any case, it is always better to slightly overestimate uncertainty than underestimate it.

6. Conclusion

Allow me to be blunt: amateur astronomers need to do a better job of calculating and plotting uncertainty. It used to be, after a long night of acquiring data, that the first thing I did in the morning was plot the light curve. Now that is the second thing I do; the first thing is calculate my uncertainty. This gives me confidence in my light curve, or in the worst cases, lowers my expectations of the quality I am going to see. The skill of a photometrist is directly proportional to the precision of his or her work. You can’t improve your skills if you don’t know how good or bad they are. In addition to measuring variable stars you can also measure your skill at measuring variable stars. Far from being a chore, it is a delight. Uncertainty analysis can be a very rewarding part of photometry.

7. Acknowledgements

With kind thanks to Arne Henden for years of patience during my constant badgering about these sorts of topics and for reviewing this manuscript. Also thanks to Michael Richmond for helpful comments and suggestions. Finally, special thanks to the AAVSO and AAVSO members for putting up with 10,000 mail list messages from me over the years.

8. References


The Magnitude and Constancy of Second-Order Extinction at a Low-Altitude Observatory Site

Bob Buchheim
Lockheed Martin Corp.
Altimira Observatory
rbuchheim@earthlink.net

Abstract
Second-order extinction is routinely assumed to be small, nearly constant, and only significant for B band or (B-V) color index measurements. This study addressed the question, "are these assumptions valid at a typical low-altitude amateur observatory site?" Happily, they are. This result should give comfort to other photometrists who use sites that are not located on high mountain tops. © 2005 Society for Astronomical Science.

1. Purpose
For many projects, determination of atmospheric extinction is necessary in order to report the standard (exoastronomical) photometry of a target object. First-order extinction must normally be determined each night; but there are efficient methods for doing this with a small investment of the night's observing time. The determination of second-order extinction requires a special set of observations, which (if done every night) would use up a significant amount of observing time. Therefore, many photometrists rely on the conventional wisdom that "second-order extinction is small (typically \( k_{bv}'' \approx 0.04 \)), it does not vary from night to night, and it is negligible for v and r band." This study of second order extinction had two purposes:

First, the conventional wisdom is validated by experience from professional observatories situated at high altitude sites. But: is second-order extinction small, and is it invariant, when observing at a typical amateur observatory site? Altimira Observatory conducts photometric studies from a location in southern California, near the coastal plain, at only 183 meters ASL. Hence, it seemed worthwhile to study the magnitude and night-to-night variability of second-order extinction.

This aspect of the study was encouraged by Landolt's (ref 1) observation that even under the nearly-ideal conditions of the Cerro Tololo Inter-American Observatory, there is a measurable range of variation in second-order extinction coefficients.

Second, the standard definitions of second-order extinction are based on observations in B and V bands. But with modern CCD’s, it is often more convenient to observe in V and R bands. Therefore, an extension of the standard second-order extinction equations to consider observations made primarily in V and R bands has been developed and investigated.

2. Relevant Equations
Following Henden and Kaitchuck (ref 2), and using their nomenclature, the observed v-magnitude of a target is:

\[
\begin{align*}
v &= v_0 + k'_{v} X + k''_{v}(b-v)X \\
\end{align*}
\]

and the observed b-magnitude is:

\[
\begin{align*}
b &= b_0 + k'_{b} X + k''_{b}(b-v)X \\
\end{align*}
\]

Subtracting Eq. 2 from Eq. 1 gives the (b-v) color index

\[
(b-v) = (b-v)_0 + k'_{bv}X + k''_{bv}(b-v)X
\]

and shows that the first- and second-order extinction coefficients for color index are related to the extinction coefficients in the individual bands by

\[
k'_{bv} = k'_{b} - k'_{v}
\]

and
Second-Order Extinctions - Buchheim

\[ k''_{bv} = k''_b - k''_v \]  (4)

Although not often seen in photometric reports (because it is presumed to be negligible), the second-order extinction effect on \( r \)-magnitude and \( (v-r) \) color index are:

\[ r = r_0 + k'_{r} \cdot X + k''_{r} \cdot (b-v) \cdot X \]  (5)

and

\[ (v-r) = (v-r)_0 + k'_{vr} \cdot X + k''_{vr} \cdot (b-v) \cdot X \]  (6)

where

\[ k''_{vr} = k''_v - k''_r \]

In the above equations,

- \( b, v, r \) = measured instrumental magnitudes
- \( b_0, v_0, r_0 \) = exoatmospheric instrumental magnitudes
- \( X \) = air mass of the observation
- \( k' \) = first-order extinction coefficients
- \( k'' \) = second-order extinction coefficients

It is important to note that these second-order extinctions are defined relative to the \( (b-v) \) color index of the target, so that the second order extinction is measured in “magnitudes per air mass per magnitude of \( (b-v) \) color”. In Section 7, I will extend this definition to consider a second-order extinction that is anchored in the \( (v-r) \) color index.

3. Determination of Extinction

3.1 Second-order \( v \)-magnitude extinction: The standard method of determining second-order extinction is to measure a red-blue pair of stars at a wide range of air masses. Applying Eq. 1, we expect star #1 to follow the equation:

\[ v_1 = v_{0,1} + k'_{r} \cdot X + k''_{r} \cdot (b-v)_{1} \cdot X \]

and similarly for star #2:

\[ v_2 = v_{0,2} + k'_{r} \cdot X + k''_{r} \cdot (b-v)_{2} \cdot X \]

Subtracting these two equations and rearranging gives

\[ \Delta(v) = k''_{r} \cdot \Delta(b-v) \cdot X + \Delta(v)_0 \]  (7)

where

\[ \Delta(v) = v_1 - v_2 \]

and

\[ \Delta(v)_0 = v_{0,1} - v_{0,2} \]

Eq. 7 is linear: a plot of \( \Delta(v) \) vs. \( \Delta(b-v) \cdot X \) will be a straight line, whose slope is the second-order extinction, \( k''_r \).

\[ \Delta(v) \]

\[ \Delta(b-v) \cdot X \]

slope = \( k''_r \)

Determination of first-order extinction is then done by applying Eq. 1. Rearranging and grouping the terms gives:

\[ v = [k'_{r} + k''_{r} \cdot (b-v)] \cdot X + v_0 \]  (8)

Thus, a plot of \( v \) versus \( X \) for any single star will be a straight line, whose slope is given by the term in square brackets:

\[ > \]

slope = \([k'_{r} + k''_{r} \cdot (b-v)]\)

Note that the first order extinction is not simply the slope of the plot of \( v \) vs. \( X \). Eq. 8 explains something that most photometrists have probably observed in their data. If you use more than one “comp star”, and create plots of \( v \) vs. \( X \) for each comp star, the lines will have slightly different slopes. This is the signature of second-order extinction on comp stars of different colors.

3.2 Second-order extinction of \( b \)- and \( r \)-magnitude: For completeness, I note that the same method of section 3.1 can be applied to determine second-order extinction in \( b \) and \( r \). The relevant equations (based on following a red-blue pair through a wide range of air mass) are:

\[ \Delta(b) = k''_{b} \cdot \Delta(b-v) \cdot X + \Delta(b)_0 \]  (9)

\[ \Delta(r) = k''_{r} \cdot \Delta(b-v) \cdot X + \Delta(r)_0 \]  (10)

The only tricky thing to notice in Eqs. 9 and 10 is that they both use \( (b-v) \) as the color index that describes “how blue or red is the star of interest?”
Similar to Eq. 8, the plots of instrumental magnitude vs. air mass (used to determine first-order extinction) in these two bands are described by:

\[ r = [k'_r + k''_r(b-v)]X + r_0 \]

\[ b = [k'_b + k''_b(b-v)]X + b_0 \]

so again, the slope of the line of IM vs. X is not the first-order extinction – the term involving \( k'' \) and color index must be recognized.

### 3.3 Second-order extinction of \((b-v)\) color index:

Following the same procedure as above, the second-order extinction coefficient for \((b-v)\) color index is found by following a red-blue pair of stars through a range of air mass, and applying Eq. 2:

\[ \Delta(b-v) = k''_{bv} \cdot \Delta(b-v) \cdot X + \Delta(b-v)_0 \quad (11) \]

where

\[ \Delta(b-v) = [(b-v)_1 - (b-v)_2] \]

and

\[ \Delta(b-v)_0 = [(b-v)_{0,1} - (b-v)_{0,2}] \]

According to Eq. 11, a plot of \( \Delta(b-v) \) versus \( \Delta(b-v) \cdot X \) will have a slope equal to the second order extinction, \( k''_{bv} \).

First-order extinction is determined by applying Eq. 2. Rearranging and grouping the terms gives:

\[ (b-v) = [k'_{bv} + k''_{bv}(b-v)]X + (b-v)_0 \quad (12) \]

Thus, a plot of \( (b-v) \) versus X will be a straight line, whose slope is given by the term in square brackets.

As a practical matter, I find it to be more convenient to determine the second-order extinction separately for \( b \) and \( v \) (using Eq. 7 and Eq. 9), and then calculate \( k''_{bv} \) using Eq. 4.

### 4. Observed values of first- and second-order extinction

I devoted several nights to monitoring a red-blue star pair at Altimira Observatory. This data set provides some insight into both the typical values of extinction at a low-altitude observatory site, and the constancy of the extinction values.

#### 4.1 Details of 01-18-2005 UT observations and analysis:

A typical set of results, which I will present in some detail in order to describe the data analysis methods, is from UT 01-18-2005. The imaging sequence was RR-VV-BB-… and since there are a few minutes between exposures, the air mass changes slightly between filters. While it would be possible to construct an “average” color index and “average” air mass bridging adjacent color exposures, it is more convenient to analyze the data in terms of individual colors, and then derive the color-index extinctions.

Figure 1 shows the observed magnitude vs. air mass in \( b \), \( v \), and \( r \) bands for six stars. They illustrate the effects predicted by Eqs. 1, and 8:

- the curves quite accurately fit linear trend lines (typical correlation coefficients are \( R^2 \approx 0.9 \) or larger)
- each star has a unique slope (e.g. looking at the \( b \)-band curves, star #3 has a slope = 0.2447, while star #t has a slope = 0.3028).

The raw instrumental color indices of these 6 stars, as measured at air mass \( \approx 1.2 \), are shown in the table below. Star “t” is HD 50279, and star “1” is HD 50167. This pair is one of the recommended “red-blue” pairs for determining second-order extinction. Note that stars 2 and 3 also provide a wide range of color, and so provide a second pair for second-order extinction determination, in the same CCD field of view.

<table>
<thead>
<tr>
<th>star#</th>
<th>t</th>
<th>1</th>
<th>2</th>
<th>3</th>
<th>4</th>
<th>5</th>
</tr>
</thead>
<tbody>
<tr>
<td>(b-v)</td>
<td>0.63</td>
<td>1.78</td>
<td>0.59</td>
<td>1.85</td>
<td>1.55</td>
<td>0.91</td>
</tr>
</tbody>
</table>

The star pairs (t, 1) and (2, 3) were used for second-order extinction determination in \( b \), \( v \), \( r \) bands. The relevant data is plotted in Figure 2. The resulting determinations of second-order and first-order extinction are described in the next section.
Figure 1: b, r, and v 1\textsuperscript{st} order extinction plots for UT 01-18-2005
4.2 Extinction Results on 01-18-2005 UT: The data from Figure 2 is used (with Eq. 4) to calculate the second-order extinctions. The results are shown in Table 1:

<table>
<thead>
<tr>
<th></th>
<th>k''_b</th>
<th>k''_v</th>
<th>k''_r</th>
<th>k''_bv</th>
<th>k''_rv</th>
</tr>
</thead>
<tbody>
<tr>
<td>using stars (t,1)</td>
<td>-0.03</td>
<td>0.01</td>
<td>0.01</td>
<td>-0.04</td>
<td>0.00</td>
</tr>
<tr>
<td>using stars (2,3)</td>
<td>-0.04</td>
<td>0.00</td>
<td>0.00</td>
<td>-0.04</td>
<td>0.00</td>
</tr>
<tr>
<td>average</td>
<td>-0.03</td>
<td>0.00</td>
<td>0.00</td>
<td>-0.04</td>
<td>0.00</td>
</tr>
</tbody>
</table>

Table 1: Second-order extinctions (from Figure 2 data)

Note that the magnitude of second-order extinction for b-band is much larger than any of the others, and that the v- and r-band second-order extinctions are negligible (consistent with the conventional wisdom). Also, it doesn’t matter much whether star pair (t,1) or star pair (2,3) is used – the measured values for k''_b (and k''_bv) are similar, and the other values are essentially equal to zero.

In the following section, the results from other nights (calculated in the same way as described here) are combined in a statistical analysis of the night-to-night variation of the second-order extinction coefficients.
5. Night-to-Night Constancy of Second-order Extinction Coefficients

Determination of k’ and k” on several nights was done to address a few fundamental questions:

- what are “typical” values of second-order extinction for my (low-altitude) observatory site?
- is second-order extinction (approximately) constant from night-to-night?
- does second-order extinction correlate with first-order extinction?

Second-order extinction was determined on 7 nights that appeared subjectively to be “good” nights for photometry – clear, stable, and acceptable seeing.

Figure 3 shows the range of second-order extinction results that were observed.

The second-order extinctions in v and r, and (v-r) are close enough to zero to be negligible for most purposes, in comparison to other error sources. Second-order extinction in b and (b-v) is significant, and shows a modest level of variation from night to night – large enough that use of an “average” value will generate a couple hundredths of a magnitude uncertainty in reported photometry – but using an “average” value is better than ignoring the effect altogether!

There is a hint of a weak correlation between first- and second-order extinction in b-band (i.e. the second-order extinction k”\( {b} \) is more negative on nights when first-order extinction k’\( {b} \) is larger), as shown in Figure 4.

This possible correlation needs more data, but if confirmed may offer a way to improve photometric accuracy when using an “average” (i.e. average trend line) value of k”\( {b} \).
6. A way of checking 2nd order extinction using comp star data in the target field

Asteroid light-curve projects require monitoring the target continuously for the entire night. Determination of total extinction \[ k' + k''(b-v) \] is thus a “free” benefit – the comp stars are being followed through a large range of air mass, so Eq. 8 can be applied.

Look again at Eq. 8, and the little graph sketched below it. These offer a way to decide if second-order extinction is significantly affecting the night’s observations with the selected comp stars. The idea is to use the fact that each star has a slightly different slope on the IM vs. X curve. Call that slope \( m \). According to Eq. 8,

\[
m = k'_v + k''_v(b-v)
\]

that is, each star gives its own value of \( m \), and if we plot \( m \) versus \( (b-v) \), we expect to get a linear graph, whose slope is the second-order extinction, and whose y-intercept is the first-order extinction. With modern spreadsheets, that calculation is a trivial exercise.

For example, using the data given in Figure 1, recognizing that each star has slightly different slope, and using the \( (b-v) \) color index of each star (measured from the same images that give the IM vs. X curves), we get Figure 5.

That is, this method gives \( k'_b = .31 \) mag/air mass, and \( k''_{bv} = k''_b - k''_v = -.03 \) which is virtually identical to the value obtained by the standard method of determining second-order extinction (compare with Table 1).

The reason that this may be a valuable observation, particularly for asteroid photometry, is that MPO Canopus permits the use of up to 5 comp stars. If the target field is imaged in \( b \) and \( v \), and if (by luck) the selected comp stars span a wide range of color index, then the plots of IM vs. X contain information for both first- and second-order extinction.

Warning: there is a bit of mathematical guile hidden in Eq. 12. In truth, \( (b-v) \) is not a constant; it is actually a slowly-varying function of the air mass \( X \). As a practical matter, however, using the average \( (b-v) \) over the air mass range used in a given night’s session, and treating it as though it were constant, seems to give reasonable and consistent results.

7. Second-order extinctions anchored in \((v-r)\) measurements

The term \( (b-v) \) in Eqs. 1, 3, and 5 describes the color of the star (not the band in which it is being observed). While it is standard practice to use the \( (b-v) \) color index, there is no fundamental reason that we must describe the star’s color by that particular color index. The CCD’s that many photometrists use are much more sensitive in \( v \) and \( r \) bands than they are in \( b \); therefore, it is of interest to examine a variation on the standard definition of second-order extinction that is anchored in the \( (v-r) \) color index rather than \( (b-v) \). I’ll call the first and second-order extinctions so defined as \( j'_v \) and \( j''_v \) (to avoid confusion with the conventional extinctions \( k'_v \) and \( k''_v \)). I was led to this idea by Villata et al (ref 5), who used a second-order extinction definition that is identical to Eq. 15 below.

By analogy to Eq. 1, the defining equations for these alternate second-order extinctions based on \( (v-r) \) are:

\[
v = v_0 + j'_v X + j''_v (v-r) X
\]

\[
b = b_0 + j'_b X + j''_b (v-r) X
\]

\[
r = r_0 + j'_r X + j''_r (v-r) X
\]

The second-order extinction coefficients are measured as before following a star-pair through a wide
range of air masses, applying an equation analogous to Eq. 7:

\[ \Delta(v) = j''_v \Delta(v-r) X + \Delta(v)_0 \]  

(16)

and constructing the indicated linear plot:

![Linear plot](image)

Several questions were investigated for these “alternate” second-order extinctions that are anchored in (v-r) as the defining color index:

- can the \( j'' \) be effectively determined? are they consistent?
- are the \( j'' \) significantly different from \( k'' \)? Is the difference understandable?
- are the first-order extinctions \( j' \approx k' \) (i.e. is first-order extinction unaffected by the use of \( j'' \) vs. \( k'' \)?)
- are target-object exoatmospheric magnitudes determined by use of \( j'' \) and \( (v-r) \) tolerably identical to exoatmospheric magnitudes determined by the “standard” parameters \( k' \), \( k'' \), and \( (b-v) \)?

The answer to all of these questions turned out to be positive.

The observed range of second-order extinctions, using (v-r) color index to describe the star’s color, is shown in Figure 6:

![Figure 6](image)

Again, second-order extinctions \( j''_b \) and \( j''_{bv} \) are negative, and large. This is only of academic interest, since as a practical matter if you’re using (v-r) as the reference index, it’s probably because you aren’t observing through a “b” filter.

The values of \( j''_v \), \( j''_r \), and \( j''_{vr} \) may be surprising at first – they seem large enough that it may be wise to account for them in color-index determinations. The reason that they appear to be (relatively) large is that the range of (v-r) is smaller than the range of (b-v) for main-sequence stars. The instrumental color-color diagram (uncorrected for extinction) for the stars used for most of these extinction measurements is very closely matched by:

\[ (v-r) = 0.69 \ (b-v) + 0.13 \]

so that it isn’t surprising that the second-order extinction defined using (v-r) is a factor of \( 1/0.69 = 1.4 \) larger than second-order extinction defined using (b-v).

Considering the difficulty of achieving accuracy better than ±0.02 magnitude in most projects, it seems warranted to continue following the conventional wisdom that “second order extinction in v, r, and (v-r) is negligible”.

8. Comparison of Altimira Observatory with other sites reporting extinction values

A search for references reporting extinction measurements at professional observatories gave an interestingly wide range of representative extinction values, which are summarized in Table 2:
Table 2: Representative Extinction values reported by professional observatory sites

<table>
<thead>
<tr>
<th>Site</th>
<th>k',</th>
<th>k'_bv</th>
<th>k'',</th>
<th>k''_bv</th>
<th>comments</th>
<th>ref</th>
</tr>
</thead>
<tbody>
<tr>
<td>La Silla</td>
<td>0.00</td>
<td>-0.035</td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Bochum</td>
<td>0.130±0.014</td>
<td>0.120±0.013</td>
<td>0.00</td>
<td>-0.014</td>
<td></td>
<td></td>
</tr>
<tr>
<td>ESO 50 cm</td>
<td>0.119±0.040</td>
<td>0.134±0.010</td>
<td>0.00</td>
<td>-0.028</td>
<td></td>
<td></td>
</tr>
<tr>
<td>CTIO</td>
<td>0.140</td>
<td>0.090</td>
<td>-0.25</td>
<td></td>
<td>affected by El Chincon eruption</td>
<td>4</td>
</tr>
<tr>
<td>SPO</td>
<td>0.250</td>
<td>0.100</td>
<td>-0.33</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>UAO</td>
<td>0.250</td>
<td>0.060</td>
<td>-0.33</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Cerro Tololo</td>
<td>0.152</td>
<td>0.124</td>
<td>-0.23</td>
<td></td>
<td>average (13 year period)</td>
<td>1</td>
</tr>
<tr>
<td>Cerro Tololo</td>
<td>0.099 - 0.25</td>
<td>0.074 - 0.184</td>
<td>-0.046 to +0.013</td>
<td>total range (13 yr period)</td>
<td>1</td>
<td></td>
</tr>
<tr>
<td>VBO</td>
<td>0.224</td>
<td>0.193</td>
<td>0.01</td>
<td>+0.026 (sic)</td>
<td>elev = 725 m ASL</td>
<td>6</td>
</tr>
<tr>
<td>Cesco</td>
<td>0.160±0.055</td>
<td>0.141±0.024</td>
<td>0±0.009</td>
<td>-0.032 ± 0.006</td>
<td>elev = 2500 m ASL</td>
<td>7</td>
</tr>
</tbody>
</table>

ESO = European Southern Observatory
CTIO = Cerro Tololo International Observatory
SPO = Sacramento Peak Cloudcroft site
UAO = University of Arizona Mt. Lemon
VBO = Varnu Bappu Observatory (Kavalu, India)
Cesco = Estacion Astronomica “Dr. Carlos Ulrrico Cesco”, of Felix Aquilar Obs., San Juan, Argentia.

9. Acknowledgements:

Photometric reductions were done using Brian Warner’s MPO Canopus/PhotoRed program. Charting and star identifications were done with Chris Marriott’s SkyMapPro. Automated observatory control was done using the Software Bisque suite: TheSky, Automadome, and CCDSoft.

Sections 2 and 3 of this paper draw liberally from Chapter 4 of Henden and Kaitchuck (ref 2). I used their equations and symbology for the convenience of readers who may want to cross-reference with ref 2. I have also explicitly stated some equations that they left as “exercises for the reader” because the subject of second order extinction is pretty confusing for some of us.

10. References


Asteroid Photometry: Tricky Business

John Menke
Menke Scientific Ltd
22500 Old 100 Rd
Barnesville, MD 20838
John@menkescientific.com

Abstract
Two nagging issues (among others) that tend to afflict asteroid photometrists (and others) are (1) concerns about just what a Flat is doing (and whether it makes things better or worse), and (2) how to handle those pesky stars that keep jumping into the path of the asteroid as it cruises across the field of view (FOV) using StarZap, a program written to subtract the offending stars. We’ll discuss experiments done to get a handle on these two issues. © 2005 Society for Astronomical Science.

1. The Flats Experiments

Many of us have trouble figuring out just what flats are measuring, and what they do, and why they work or don’t work. This is NOT a talk about T-shirts vs. sky flats (though I will touch on that). It is a talk about some of the more subtle aspects of understanding just what flats are measuring, and when they may be misleading.

All of us have done flats of some sort at some time. My own first experience was when I tried imaging with a x3 Reducer on a 6 in. f/12 AP refractor. The vignetting was obvious, with a loss of some 30% of intensity toward the edges of the field (and this was with what is now considered a small field ST7). I tried doing a flat field, and lo and behold, the vignetting disappeared, leaving only a little extra noise in its wake. I was having little problem with dust in my camera, so vignetting was the only major problem. I tried different flat techniques (sky, Tshirt (or sheet over the scope), dome flat with a Styrofoam panel) and all seemed to give the same result. Of course, I kept reading the religious wars over flat techniques on the SBIG list and elsewhere, but it didn’t seem to apply to me. But I also kept having this nagging thought—why does a dome flat from six feet away give the same result as a sky flat from five miles? And is it really the same?

My ignorance continued with no challenge until I started doing photometry, first of stars, then of asteroids. Some parts of the photometry were easy, some very confusing. But when I tried doing flats, I saw little effect (I was back to using an f/6 reducer on the C1, ST7E, with very little vignetting), so I often skipped them. But then my V-filter got dirty. Of course, I didn’t know it at the time, as I was doing remote observing and didn’t look at the equipment for weeks or months at a time. But image quality started deteriorating, and funny things started happening. For example, I would flip the GEM mount, and the star and references would show a changed relative brightness offset of perhaps .05 mag as they landed on different parts of the field.

Well, I thought, that just means that the effective sensitivity across the FOV is not constant, so I’ll just do a flat to get rid of the effect. Well, that helped—sometimes. But other times the flat seemed to make things worse. So I began trying to think about flats, and what they measure, and what they don’t. I must tell you that while I have made some headway in both thinking and experimenting, I do not have all the answers nor do I fully understand what is going on. But I will give you a hint of what is to come: I was baffled.

One of the first questions I tried to understand was what problems a flat can help. The purpose of the flat is to compensate for sensitivity variations across the FOV. You can have a sensitivity variation in the CCD itself (though I’ve never seen this), or it can be introduced by any variations in the light path from source to chip that are different from one part of the chip to another. There are two general effects we will discuss (1) dust doughnuts, (2) gradients or gradual sensitivity variations across the field.

Let’s take the easiest case of something that will cause a variation of sensitivity at different parts of the chip—a speck of dust on the chip. But even that is not so simple: the speck is really not on the chip, it is probably on the cover slip that is on the chip. Furthermore, the speck is not opaque, so has some variable transmission. Let’s look at a typical geometry for an ST7 on an F/6 scope as shown in Fig. 1.
We talk about a “light cone”, but really, every pixel is receiving its own light cone. But we can see that the speck will cast a partial shadow on the CCD as it intercepts each light cone, and the pattern of the shadows on the CCD will be a doughnut. The radius of the donut depends on how far away the speck is from the CCD, and the convergence (i.e., f number) of the light cone(s). You can measure the radius in pixels, multiply by pixel size and f number, and compute how far from the chip the dust is. The % of darkening or obscuration of the dust will depend primarily on the size of the speck, and on the f/number and so on. In general, the cone of light for one pixel is much, much larger than the dust speck where it crosses the speck so that even an opaque speck will cause only a partial blockage (attenuation) of the light.

We can actually visualize this process by putting a cover on the scope that is pierced by small (1/8”) holes. For example, a single hole will create a very narrow cone of light that will project the shadow of the speck onto the chip (this forms, in effect, the light cone from an f/900 scope). More holes, even spaced, will show how the pattern builds into the familiar doughnut. Some of these test results are shown in Fig. 2. On the right is an image through one hole, on the left is through 18 holes. Geometry shows that these dust doughnuts come from the CCD window.

In most astronomy optical systems, what you will probably find is that the smaller dust doughnuts you see (the intense, little guys with a diameter of 5% or less of the FOV) are inside the camera on the cover slip. Dust on the camera window will show much larger doughnuts (e.g. diameters 10-30% of the FOV). Dust on the filters will form even larger, but fainter, doughnuts. Clearly, once you leave the region close to the chip, dust has a rapidly decreasing effect in terms of obscuration: most of the light just goes around it to get to each pixel.

But the question still remains: how do we correct for the attenuation, pixel by pixel? If we can measure that attenuation pixel by pixel, then we can correct the measured intensity at that pixel, i.e., do a flat calibration.

So how do we do that? In theory, we could put a parallel beam into the scope that creates a light cone the size of a single pixel. We could measure the sensitivity of that pixel, then move on to the next, and map out the whole system. This is probably the ideal way to measure a “flat”: we’ll come back to why that is so. But that is slow and difficult. So, instead of doing one pixel at a time, we do them all at once (so they each calibrate one another: We flood the chip with uniform light (at least, that is what we TRY to do), and measure the relative sensitivity of all the pixels at once.
The amazing thing is that this works so well: it is truly a quick and dirty technique. But with a little thought you can see immediately several problems with this approach. The major concern is that the flat is taken by flooding the field and these are VERY different conditions from how you actually take an image. Differences include:

- Total amount of light hitting the CCD is different (usually much, much larger) and in a different pattern which may cause variations in response across the CCD (e.g., either from chip non-uniformity, or from electronic non-linearities as the electronics copes with the high readout values)

- With light flooding the field, the light reaching the speck affected pixel may be augmented by light scattered within the camera, or elsewhere in the optical system. This will affect the flat calibration accuracy.

- The scattered light reaching the camera may depend on the range of angles of light entering the telescope, the baffling of the telescope, and any field stops in the optical system. The range of angles of light rays entering the scope may be very different between 'Tshirts, dome flats, and sky flats. For example, the mix of light ray angles generated by imaging –i.e., in-focus- a 2 foot white circle at 500 feet away is not the same vs. a way out-of-focus T-shirt on front of the scope or a sky flat.

- The flat light may be a different color from the image light

Then there are always statistics: the flats you take always have statistical errors, so there will always be at least some adverse effect on the noise of the flat calibrated images.

So, how come flats work at all? Well, for imaging, the demands are not too high. If you can get rid of doughnuts and obvious vignetting, you will likely be satisfied with the resulting image. After all, if there is a decrease in brightness in one part of the image, who is to know that it is not in the object? Also, the intensity range across most images is in the 100:1 range, so a few percent of smooth variation will not be apparent.

The situation, of course, can be very different for photometry. An error in the flat between the target and the reference star(s) pixel locations may cause no problem if the target and reference stays on the same pixels all night (or if there is a uniform gradient in the direction of their motion)—e.g., the target and reference will always be X.XXX mag different. However, if the gradient is not uniform, or if the target and reference stars move to different pixels (imperfect tracking, a GEM flip which may even interchange target and reference, movement of an asteroid) then an error will have been introduced by the flat calibration. You may detect that an error is present, but it may be very difficult to correct for it.

So, what to do? How can you know how good your flats really are?

The easy answer is to take your photometry data and treat it with and without flat calibration—and hope the answers are the same, thus probably showing that you are ok. You watch carefully for signs of error—again, both with and without flats. You work with a system that is as clean as you can (to reduce doughnuts), and that has good baffling to reduce stray light.

The hard answer is that you can do experiments similar to what I will describe. Over the past several years, I have done the following sets of tests to try to understand what is happening:

- Compare Tshirts (or sheet on scope), dome flats, and sky flats
- Perform field tests in which I presented a well defined cylinder of light (first from a 50 gal drum, then a much fancier octagonal light source—which did no better) into the scope, i.e., control the angular spread of incoming light. I can compare this at very close distance (e.g., 20 ft) with long distance (120yd). Making a 30 inch diameter flat source has been an interesting project, but one that so far has not reached the 1% flatness I was seeking.
- Perform basement tests using sheets and lights in various configurations (distances up to 50 ft) both with the camera alone, and with the camera on a variety of telescopes
- Tests using true star fields to evaluate the uniformity of the CCD response, and to compare to flats.

Virtually all outdoor tests were performed with the ST7E, CFW8, on a C11 operated at f/6. I obtained and analyzed the images using MaximDL, which provides easy to use graphing and analysis tools.

In any test, there is always the question of whether a given artifact (e.g., a gradient across the field) arises from the flat source (or its lighting), the scope, or the camera. To try to identify the source of an asymmetric artifact, one can rotate each of the components in the system and try to track the change in the artifact. Keeping straight what you are doing is a real challenge (and even harder, is reconstructing what you did six months later). You may have the camera, filter, scope, scope hole pattern, flat, and flat lighting to rotate. Is that six test alternatives, or \(2^6=64\), or only 32. Well, it depends. But regardless,
it is a pain because rotating some of these items is easier said than done. Also, errors that are symmetric about the optical axis cannot be teased out just using rotations.

As I noted, I had compared dome flats, Tshirt flats, and sky flats. They produced similar results, but they had disturbing variations under different conditions which I could not identify. For example, sometimes there were 4% gradients across the FOV, other times is was only 2%. While none of these are terrible for .05mag photometry, they did show that I did not understand my system.

So I stripped the system to the essentials. Working in my basement, I set up the experiment shown in Fig. 3. I aimed the camera through a 3" D by 36" tube fitted with field stops as shown. This was aimed at a white sheet five feet away, which was illuminated by a light as shown. I could put a diffuser (piece of paper) at A (on the camera nosepiece) or at B (end of the tube). (I also did other experiments, but will only report on this set).

![Figure 3 Diffuse Light Source Test Setup](image)

With a diffuser at B, we have a sort of f/36 light source: the largest angles of the light with respect to the axis are as if it were f/36. With such simple geometry, I should be able to understand what is going on.

Figure 4 shows some of the results. The top pair 4A&B are for the unmodified camera. In this and following figures, I have normalized each image to its total brightness so that we can easily compare sensitivities. Note that a brightness change from 1.01 to 1.02 is a 1% (.01 mag) difference. Also remember that positive Y is downward. The Fnn designation is my own nomenclature for my images (not f/number).

Fig 4A (upper left) shows a terrible vertical gradient of more than 5%, with Fig. 4B showing a substantial gradient across the middle of about 1.5%. This seemed terrible! What is going on? Remember, this is an f/36 like system—the "shadow" in Fig 4A is NOT the familiar case of the guide chip mirror shadowing the image which can produce a very narrow dark band at the top of an image taken with a fast optical system. I had also seen similar behavior as this in my flats taken with the C11 at f/6 and even f/10.

After much experiment, thought and false steps, I postulated that it might be due to scattered light entering the chip. I installed a simple mask, cut large so as not to cut off f/6 light rays. The result is shown in Figure 4C—a major change! The vertical gradient virtually goes away, and the horizontal gradient is now much more regular.

Clearly, baffling is important, and this is proof that at least in my ST7 there can be a major problem with scattered light. What is it doing? It is entering the camera, bouncing around, and changing the pattern of light on the chip itself. That is, the scattered light contributes to the flat. And what is wrong with that? What is wrong is that it causes the flat calibra-
tion process to produce errors in the corrected images! This is because the stars we usually measure are superimposed on a very weak background of similarly scattered light which causes no problem because we are doing aperture photometry. But the flat calibration introduces a calibration variation that varies across the FOV, and which does not in fact represent actual, varying sensitivity. Which is what I had suspected.

As I looked at Fig. 4C, I realized I now did not understand the bell curve: was it the mask cutting off the edges of the light cones? But no—the mask should not have been doing that in this f/36 system. Again, after more thought, I installed small stops in the camera nosepiece which gave Fig 4D. The bell went away, and the curve is flat to within 0.5%. Clearly, some of the f/36 rays were reflecting off the interior of the nosepiece or part of the camera into the central area of the FOV.

The basement tests (which included test with several telescopes, and use of a diffuser at A) helped me understand flat processes, and showed me that I was having a real problem with scattered light which the mask helped. Obviously, in a well baffled, high f-number system, these effects will be small. But how could I actually prove how good my system in the observatory really is?

I then did star tests. Remember, I said it would be ideal to run tests on each pixel, one at a time? Well, I didn’t do that, but I could do a subset of those tests using stars.

One way to do this would be to take an image of a star field (using the C11 at f/6, mask on camera), then translate and/or rotate the scope/camera and evaluate the measurements to determine the spatial pattern of response across the image. Doing this thoroughly for even 10-20 stars would be a lot of work, so I decided to try just two stars. I picked two stars of similar brightness about 5 a-min apart on an E-W line (there aren’t very many!). My FOV is about 12 a-m wide. I could thus center StarA image, then center StarB. When centered, A & B in the two images are at the same spot on the chip, so have the same calibration which we use as the reference. We can thus flick the scope back and forth and build a sequence of images. We can measure the relative brightness of the star pair Left, then Right, to see whether there is a consistent difference. While this won’t detect a constant gradient, it will identify a symmetric error or non constant gradients. I then rotated the scope and camera 180deg, and repeated the test. I then rotated the camera on the already rotated scope, and retested.

Even using multiple images, the scatter in the data limited the accuracy of the measure of gradients to about 0.5%. However, the data did show that there was less than about this much (0.5%) variation across the image. Great! Although the measurements could not rule out that there are still some issues of scattered light, there is clearly no systematic problem of calibration, at least in the horizontal band studied.

I then did a final series of flat tests including a Tee shirt flat through my Red filter of the daytime sky, a night-time sky flat (using local light pollution), and sky flats with the tube/camera rotated, then only the camera rotated back.

When I plotted the results, I found that the curves are not as nice (smooth) as those from the basement tests although they were generally consistent. The deviations showed I still have some scattered light entering the chip, but far less than when I started.

Because the star measurements showed less than 0.5% gradient errors, if I used typical flats with 0.5-1% variations, I would apparently be introducing errors! However, it is also clear that all these effects contribute less than .01mag error across the chip (now that I have the mask on), so I can ignore them. And, of course, using the flats would remove doughnuts that might be more than 0.5% errors.

I may or may not use a flat, but at least I DO know what some of the limits on the system are! And I think I understand doughnuts.

Flat Summary

At least for my camera, scattered light can cause systematic errors in flats, and introduce systematic errors into your data if you use those flats in calibration. Use of wide angle flat sources (sky flats, Tee shirt flats, perhaps dome flats) in fast systems are likely to be at greatest risk of this effect. Using star measurements, you can verify at least to a limited degree the actual performance of your system. In faster systems, you may not notice the dust doughnuts on the flats taken through filters, so you should inspect and clean the filters regularly.

2. The StarZap Experiments

The major problem that makes asteroid photometry more difficult than stellar photometry is that the asteroid moves through the FOV. Depending on the distance and orbit of the asteroid, it may move only a few a-min in a night, or it may move 20-30 a-min (or much more, if it is a near earth asteroid). In any case, the asteroid is almost always in a different field every night. This requires different reference stars every night, and creates the problem of making accurate night to night matchups of the data.

But the other problem from the moving asteroid is that it, in its journey across the FOV, it seems to seek out every possible star. As the asteroid moves
across, or near, a star, the background photometry annulus will first cross the star, causing an apparent decrease in the asteroid brightness (background is subtracted from the central photometry annulus). Further on, the central annulus covering the asteroid may cross the star, thus increasing the apparent brightness. Then it decreases again as the asteroid leaves the star, looking for its next victim.

Naturally, Murphy’s Law requires that these brightness variations always fall at the points on the photometry curve that are most critical (e.g., the missing section of the curve, where it matches to the previous night, where you think there was evidence of a satellite, etc.). Obviously, the easy thing to do is to throw out the data near the star, leaving a data gap. However, this loses valuable data. And it seemed to me that it should be possible to correct for the problem at least some of the time.

The basic idea is to subtract an appropriate star image from the combined asteroid-star image, leaving the asteroid by itself (and to do this for all the affected images in the sequence). There are two obvious sources of subtracting star images:

- The subject star taken from a different image in the sequence in which the asteroid is not involved
- Some other star in the image of concern, suitably normalized to match the subject star which is mixed with the asteroid.

Of course there are pros and cons of each method. The first method is in some ways easier to implement, and in some way “better” in that it uses the image of the subject star (especially important if the subject is a double star or has a nearby star). On the other hand, variations in transparency during the sequence of images will require some sort of normalization be applied to the subtracting image. However, experiments with both approaches seem to show that the second method, to use a star image from the subject image, appears preferable. This is because the largest contributor to error in the process (other than limited statistics) is the variations in seeing and guiding between images. That is, on average it is better to use a different subtracting star (appropriately normalized) from the same image than the same star from a different image. If the offending star is in fact a close double, rather than an isolated star, then in that case it would likely be better for the subtraction to use a copy of the offending star (suitably normalized) as taken from another image. This is easy to arrange.

Using first manual experiments, then a VB program, I have experimented with this general scheme using MaximDL. Maxim is suitable because it contains many pixel math functions that can be accessed by scripting or VB programs. Other programs may also be suitable hosts for this technique.

Experiments showed that if I use MaximDL to align a sequence (e.g., 50 images), then measure the variation in where the centroids of the resulting star images are, I find that the variation tends to be in the 0.03 pixel range as measured by the variation in the x/y positions of the star centroids. Thus, I can automatically align all the images at the beginning of the process, identify stars of interest (i.e., that the asteroid comes close to) by their x/y coordinates one image, and use that information throughout the analysis.

All but one of the methods this program script uses are standard in MaximDL. The exception is the current MaximDL automatic star alignment method will not work with only one star in each image. This feature is needed to align the ref star to be subtracted to the exact position (< 1 pixel) of the target star determined from the REF image. At my request, Doug George had Hilderic Browne modify the system to permit one-star automatic alignment.

Images used in this test were taken using the ST7E at f6 on the C11 giving about 1 a-s per pixel. The setup is mounted on an AP1200 operated without PEC and without guiding over the 180 sec exposures. The average P-P tracking error is about 3 a-s on a 7 min. period, so there can be 1-2 a-s of image movement during an exposure. The FWHM of the star images is typically in the 3.7-4.3 pixel range (worse if poorly focused). I normally use a photometry aperture radius of 8 pixels, with an additional 3 pixel radius for the background annulus.

Figure 5 shows a typical sequence as the asteroid moves through a rather crowded field. As you can see, the asteroid comes very close to several stars in its path, one of which is nearly 10x brighter than the asteroid itself.

If we want the subtraction to be good to .03 mag (3%) then you need about 0.5% precision in
your subtraction—a pretty tall order as it implies .005 mag precision in the subtraction process. In some ways, the situation is worse, because if the reference star has a different profile (PSF) from the star, even if the total brightness subtracted is correct, there may be substantial under-subtraction (residuals) or over-subtractions in different parts of the star/asteroid image. On the other hand, the closer the asteroid path to the center of the star, the more these problems will tend to average out as the photometry aperture measures the total brightness over a relatively large area.

In a normal photometric analysis, we load the sequence into MaximDL, do dark subtract and perhaps flat field, and then measure the brightness of the asteroid and selected reference stars (I usually use four or five) in each image of the sequence. I export the raw data into Excel, then visually inspect the reference stars for variability, then usually choose just one as the reference star (usually one of the brighter ones). The images comprising Figure 5 yields the upper curve in Figure 6. The effects of the interfering stars are obvious.

To demonstrate how the StarZap works, let’s go back to the image sequence. As noted, I first align the sequence. I then create a summed image that shows the path of the asteroid among the stars as in Figure 5. Using this image, I note which stars are close to or on the asteroid path and identify these target star(s) by their locations (X,Y) on the image (I do this by just clicking on the stars in the StarZap program). I also identify an appropriate reference star I will be using for subtraction. This reference star will be a reasonably bright star that I can use both to measure the relative sky transparency appropriate to each image in the sequence, and as a good quality (good statistics) star image for use in the subtraction process. The program uses the reference image to measure the brightness of each subject star relative to the chosen reference star.

After choosing the stars, I then start the program which follows the steps shown in the flow chart below. The whole process takes about five minutes total.

Once the program finishes subtracting the various subject stars, I have a sequence of images that are the same as the original sequence except that they have been aligned (which actually makes the photometric analysis go faster) and have had the interfering stars removed. Once complete, I perform the usual photometry on the modified image sequence that has had stars removed.

![Asteroid LightCurves](image)

Figure 6. Asteroid Lightcurves.
How good a job does this do? Figure 6 also shows the resulting light curve for the sequence. The improvement in the light curve compared to the original curve is obvious. The apparent errors remaining in the curve after the subtraction are in the range of <0.1 mag (<10%) even for the 1000% error from the brightest star—a very substantial reduction. However, to evaluate the precise errors will require extensive comparisons of data, which I have not yet completed.

**StarZap Summary**

At this time, I have only just begun using the program as a standard tool in my data reduction so it has not yet been tested on a wide variety of qualities of data. However, because the results presented here involved NO special settings or fiddling with parameters and yet still produced such good results, it would seem this approach has the potential greatly to improve our ability to measure light curves in crowded star fields. Anyone who wishes to experiment with this program or further develop it is welcome to contact me.

### 3. Acknowledgements

I would like to thank Doug George for his willingness to modify MaximDL to support this experiment. As always, I also thank Brian Warner for his incredible encouragement of fellow asteroid photometrists. I would also thank my son Greg for his patience and commiseration as I rant about scripting and VB, and for his help in learning how to overcome the hurdles.
Spectroscopic Monitoring of Be type Stars

Valérie Desnoux
Astronomical Ring for Astronomy Spectroscopy
66 rue du theater – 75015 Paris - FRANCE
Valerie.Desnoux@free.fr

Christian Buil
Astronomical Ring for Astronomy Spectroscopy
6 place clémence Isaure – 31400 Castanet - FRANCE
Christian.buil@wanadoo.fr

Abstract
The study of Be stars is a perfect example of spectrography that amateurs can participate in. Several types of spectroscopes can be used depending of what physical parameters are monitored: classical surveys at low-resolution or high-resolution profile study of the H-alpha line are just two examples of what can be monitored. By grouping observations on long time scales, a database of stellar spectra can be built and placed at the disposal of professionals. Over a period of 10 years, starting at the Pic du Midi observatory and continuing later on by Christian Buil, and other French amateur groups using Musicos professional spectroscope, the monitoring of Be stars produced scientific results and are now supported by professional astronomers. Some are already using the ARAS (Astronomical Ring for Amateur Spectroscopy) list to ask for specific monitoring and get additional observations from the amateur community. Spectral processing will be also be discussed, using Visual Spec and Iris freeware, from the basic calibration up to time resolution H-alpha profile evolution and Equivalent Width computation. Be Stars survey is just an example of how powerful spectroscopy can be, and can be practiced by amateurs. Amateurs are really capable of playing a key part in astrophysical studies. © 2005 Society for Astronomical Science.

1. Introduction

Spectrography is quite a different technique for amateurs. It does not really produce nice images. We do not image the splendid extension of a gas nebulae, or show the details of a planetary surface. With spectrography we penetrate into the physical intimacy of the object: temperature, density, composition, evolution of a gas disk around a star. By studying the light profile spread out per its energy we get access to a larger number of astrophysical data which are of great interest for professionals.

Spectrography requires an additional device, the spectrograph, to get into the mystery of light. It also requires some specific processing to reveal the true nature of the phenomena observed. But the results are quite enthusiastic. Most of the actual astrophysical discoveries are made by professionals with a spectrograph. More than eighty percent of observation time at professional observatory requires the use of a spectrograph. It is time now for amateurs to enter this new world and contribute to the fantastic adventure of astrophysics.

2. Be Stars

B-type stars are hot and blue stars. Their temperature varies from 10000°K to nearly 30000°K. Their mass ranges from about 3 to 20 solar mass and their luminosity from 100 to 50000 solar luminosity. Spica and Regulus are two examples of B-type stars.

Be Stars are non-supergiant B stars that at least once demonstrated Balmer lines in emission. The "e" stands then for "emission". It represents about twenty percent of the actual B-type star population. First observed Be star was G Cas by Secchi in 1867.

But not all Be stars remain Be and some B-type stars can suddenly become a Be star. What's behind this phenomenon?

To observe emission instead of absorption this means that in the line of sight more electrons are in the excited stage than at their normal energy level. When looking at a star, the atmosphere of the star becomes excited by the internal energy of the star. The de-excitation of the atoms emits photons in all the directions, so on the line of sight the overall sum shows more photons absorbed than emitted, as shown in figure 1.
Fig 1. As the atmosphere of the star is directly in the line of sight the overall balance is more absorbed photons than emitted (spread in all directions) and thus an absorption line is observed.

If there is a disk of gas close to the star, some portion of the disk does not really superimpose on the line of sight to the internal core of the star. Thus, more emitted photons are seen than the one absorbed and an emission line is then visible on the spectrum, one as shown in figure 2.

Fig 2. As the star is not really in the line of sight, more emitted photons are seen than absorbed ones, so the overall balance is a line in emission.

Be phenomenon is explained by the presence of a circumstellar disk created by episodic ejections of mass by the B star. The origin of this phenomena is however still unknown.

In a Be star spectrum, the line profile observed depends also of the inclination angle of the observer’s line of sight and the axis of the disk. This drives a large number of emission line profiles in Be star spectrum. As shown in figure 3, even a sharp emission line can be seen in the middle of the larger absorption line. The emission line itself can exhibit profile variation, like in figure 3 where two peaks are visible, indicating Doppler shift in the disk of gas surrounding the star.

Fig 3. Depending of the orientation of the disk in the line of sight, an emission line can be seen superimposed in the middle of the absorption line.

When observing a Be star spectrum over time, it is even more complicated, as evolution of the emission line profile varies with time. The observed variations can be changes in the intensity, the shape or the wavelength of the line. This can be explained as indicators of modifications in the disk structure. Monitoring and recording of such variations will then be included in the stars models developed by professionals and will contribute in the understanding of the mechanisms which sustain the phenomenon.

3. Spectrograph for Be Stars

The Balmer lines are the ones most commonly seen in emission is the Be Stars. The H-alpha line at 656.28nm is the strongest one. As this wavelength perfectly fits to the spectral response of a CCD camera, it is then the preferred line for high resolution monitoring.

Be stars are not too faint stars. The magnitude of G Cas is 2.47 for example and is thus an accessible target for the amateur telescope.

Depending of the dispersion of the spectrograph, and the limit magnitude driven by the telescope aperture, different studies can be pursued, all being of interest in the study of Be stars.

The spectral resolving power is a fundamental parameter of a spectrograph. Finer details in the spectrum are observed as the resolving power increases, like a magnifying glass more and more powerful. The mathematical expression is $R = \frac{\lambda}{\Delta \lambda}$, with $\lambda$ the working wavelength and $\Delta \lambda$ the smallest distance separating two lines of equal intensity just before one cannot distinguish one from the other one. As for example, if the observed line is Hα ($\lambda = 6563$ angstroms) and if the finest detail extends over a spectral
interval $\Delta \lambda = 0.5$ angstroms, the resolving power is equal to about $R = 13000$ (this being 6563/0.5).

One speaks of low resolution when $R$ is inferior to 1000, of medium resolution when $R$ ranges from 1000 to 10000 and of high resolution with $R$ above 10000.

**Fig 4.** LHIRE2 spectrograph sampling 0.232 Angstrom per pixel – $R = 28000$

It is important to note that a spectrograph having a high resolution is not inevitably more efficient and better than a spectrograph with a lower resolution. The high resolution is obtained by spreading the light (the spectrum) over the detector. The information dilution results in loosing detectivity, i.e. it prevents the observing of fainter stars.

A trade-off exists between the limiting magnitude and the resolving power, strongly associated to the type of program considered. The corollary is that there is not a universal spectrograph, capable of observing equally well all the sky objects.

The Be star survey has this characteristic that it is of great interest whatever the resolution adopted. So, if you own a spectrograph, it can surely be used in the survey of such type of stars.

As an example, the figure 5 shows the visible spectrum (from 4300 to 6800 angstroms) of the Gamma Cas star with $R = 1000$. The interest of this resolution is that it permits to record simultaneously a very large spectral domain and a lot of lines of interest. In this actual case, the lines H$\alpha$ (on the right) and H$\beta$ (on the left) are all together in emission phase. The ratio of their intensity provides direct information on the level of excitation of the gas in the disk. In addition, a spectrograph with a resolving power of 1000 associated to a 8-inch telescope allows the study of potentially more than thousands of Be star up to the 13th magnitude with a good signal to noise ratio.

**Fig 5.** The spectrum of Gamma Cas observed with a $R = 1000$ spectrograph and a 5-inch refractor.

An intermediate resolution of about 5000 will allow you to probe a large sample of standard B stars to detect if and when one of them may become suddenly a Be type star. The case of Delta Sco is a current spectacular example of a star transitioning from B-type to Be (see the figure 15).

**Fig 6.** One of the first spectrum showing Delta Sco as a Be star. The spectrograph of $R = 8000$ is coupled to an 8-inch telescope. (Christian Buil, May 2005).

A high resolution spectrograph will allow you to study only one line at a time, but with a lot of details. Periodic and aperiodic phenomena of a very subtle nature can be observed. Evolution on a time scale of a few hours, even less, are sometimes seen, associated with sudden ejection of matter from active areas of the star. A spectrograph resolution between 15000 and 20000 coupled to a telescope of 11-inches allows the 7th magnitude to be reached, which gives access to the study of more than a hundred stars. Many so called stable stars will no longer appear stable with such equipment. As result, there is an exciting aspect to plunging into spectroscopy. Figure 6 shows the H$\alpha$ line of Gamma Cas at the resolution of $R = 18000$. 


4. Acquisition and Processing

An image of a star spectrum is at first like any other astronomical image. Standard CCD image correction applies, the main being the dark and the bias subtraction. Correcting for flat is a little bit different. If a flat image is obtained on the sky, the image recorded is the spectrum of the sky, not a flat field. Same problem if you make a flat of a cardboard illuminated by a lamp. All you obtain is the spectrum of the lamp. So if you did not like flat fielding for classical imaging, you may not like it for spectrography. It also depends of the type of spectrograph you use, with a slit or without a slit. The main purpose of such a correction is to remove dust effects. Compensating the CCD response is usually done in another step by using a real star spectrum, well known, and not a not-well characterized lamp.

Once those corrections are done, the spectrum image is ready for more dedicated spectral corrections.

The final goal of spectral pre-processing is to convert a spectrum image into a spectral profile, a graph which displays the variation of intensity per wavelength. Very basic operation would be to simply pick a line from the spectrum image and consider this. Doing that way will throw away most of the acquired signal and data, as the spectrum image does never fills the CCD matrix entirely with only one line. The best process is to combine by addition all the lines of the spectrum to maximize the signal to noise ratio. To get into the final addition some preliminary steps are however required. First, the spectrum acquired is more or less aligned with the pixel matrix of the CCD. If adding column by column the spectrum pixel intensities, the lines will be smoothed and the resolution will be degraded. In Iris, a free software package designed by Christian Buil, dedicated spectral processing commands to correctly extract the intensity profile are present. Commands to correct for tilting, smiling or slant effect allows a fine alignment of the spectrum image to the horizontal axis of the image. Then, the spectrum of the sky background shall be removed with other commands, which allows the user to subtract the surrounding spectrum background. When recording a spectrum, the spectrum of the surrounding sky is always recorded, superimposed on the star spectrum. If not removed, lines from the sky such as the solar spectrum reflected by the moon or other atmospheric effect like telluric airglow lines will pollute the star spectrum.

The spectrum image is now ready for the final intensity profile extraction: the binning. Iris allows you to perform this final step with a dedicated command or you can load the final image into Visual Spec for an automatic binning operation. The result of the binning is an intensity per pixel graph. It is not yet a useable spectrum… the final step consist in finding the relationship between the pixels and the wavelength.

Looking at the spectrum profile, one can see lines, but it is difficult to say which lines they are, and at what wavelength until the wavelength calibration is done. Several strategies can be used.

This article will not describe all the possible ones, but will present the most robust one which led to the most accurate result. For the following spectral processing steps, the authors have used the software Visual Spec. Visual Spec is a free software, designed by Valérie Desnoux, and is dedicated to spectral analysis on a windows based PC platform.

To find the relationship between the pixels and the wavelength, a calibration lamp and the record of its spectrum in the same condition than the star spectrum is used must be taken. The calibration lamp
should have a spectrum with sharp emission lines. Here is where neon or Hg lamps are very useful. (See Figure 9)

![Neon spectrum of a calibration lamp superimposed to the star spectrum. Neon lamp lines will be used to find the relationship between the pixels and the wavelength.](image)

Fig 9. Neon spectrum of a calibration lamp superimposed to the star spectrum. Neon lamp lines will be used to find the relationship between the pixels and the wavelength.

The lines of such a spectrum can easily be identified, labeled and then used to build the equation "wavelength=f(pixel)". This relationship may not be linear, and for better accuracy, more than two lines can be used to compute the polynomial regression which will describe the relationship. Once the polynomial equation is determined, it can then be applied to the star spectrum. Each pixel will now correspond to a wavelength range which is equal to the dispersion (sampling) of the spectrum. If the spectrograph provides a limited dispersion, some uncertainty will remain on the line wavelength. As an example, on 2nm per pixel spectrum, it will not be possible to distinguish two lines which are separated by less than the sampling factor, they will be blended.

Another very interesting strategy is to use the atmospheric lines, always present on spectrum taken from earth. If you are using a spectrograph which gives you a dispersion equal or better than few tenth of angstroms such atmospheric lines are visible around the H-alpha lines, as shown in figure 10.

![Atmospheric lines around the Ha line](image)

Fig 10. Atmospheric lines around the Ha line

To improve the accuracy of the polynomial regression, such lines can be used to calibrate the spectrum. But at the final stage, they shall be removed before making any computation on the line intensity. Visual Spec provides a mean to display a synthetic atmospheric spectrum, to adjust it to the instrumental parameters of the spectrograph and to remove then by division. Thus, the spectral profile will not include any non-star perturbations.

To compare spectra over time, it is critical to record location, time of observation and to normalize the intensity. If you simply superimpose two spectrums of the same stars separated by a period of time the recorded intensities is first linked to the exposure duration on an absolute scale. So, it is recommended to look at relative intensity taking the surrounding continuum as the reference. Visual Spec allows you to define the continuum domain where an intensity mean will be first computed then will divide the entire profile by this value. The result is a relative intensity profile, relative to the continuum, which is assumed to be constant over time. (This may not always be the case, and some strong variable stars, like novae can have a strong variation of their continuum in their exploding phase, so attention should be paid). The continuum area taken as reference should be free of lines.

Another normalization process should be applied if the period of time, which separates the observations is long enough to have Doppler shift due to the earth motion. Heliocentric correction will use star coordinates and observation time to correct the Doppler shift.

5. Observation Strategy

The Be Stars survey started in France in 1990. At that time, no spectrograph was commercially available and only a small group of people in 1989 were conducting spectrographic observations. The professional observatory "Midi-Pyrénées" also well known under the name of "Pic du Midi" observatory, in the Pyrenees mountains at the border of France and Spain, had created an amateur association to use the 60cm (24 inch) telescope in 1982. In 1987, Daniel Bardin, member of the association built a spectrograph for the T60. In 1989, with the appearance of CCD cameras, the association organized a spectrography seminar. Software was DOS based. From that time, one or two weeks per year were devoted to regular monitoring of Be Stars after the discovery of the Be Atlas of AM.Hubert-Delplaces. Be stars are bright enough to allow the usage of spectrograph of R>7000 with good signal to noise ratio. This dispersion is enough to register variation of the H-alpha line. With a six month interval, some stars were already showing either intensity variations or profile shape variations. This was quite exciting, but since the Internet was not largely used, no publication of the data resulted. The T60 operation was interrupted.
in 1995. During that time, the Visual Spec software was created and the necessity to design a "home-spectrograph" was obvious. After several attempts, a "wood and metal" spectrograph, which could be used with standard amateur telescopes, was designed by Christian Buil. He re-started the regular monitoring of Be stars. We decided to combine the T60 data and his personal observations and made it available on a web site. The survey is still in progress, the Be Stars are at least observed once a year and every month or week if a rapid variation is discovered or if professionals indicate a target of interest.

As variations are not predictable, Be stars have no minimum period of time between two observations. Every spectrum counts. The more dispersive the spectrograph is, the more likely variations will be recorded.

With standard off-the shelf telescopes, with a spectrograph of R>7000, a regular monthly observation seems to be well indicated.

Another association "Astroqueyras" which exploit a 60cm telescope in the Alpes, with the observatory of "Paris-Meudon were given a professional spectrograph. The Musicos spectrograph is an "echelle" spectrograph and provides spectrum at R=35000 on the wavelength range from 4000 to 10000 angströms (fig11). In 2003, a regular campaign on Be Stars started with the goal to focus on high resolution, rapid variations of specific stars listed by the professionals or found of interest in articles published on the internet, like in the Be stars Newsletter.

There is still an area where no regular survey is happening: The B-type to Be star transitioning. As explained, the Be phenomenon is not a "once for all" star characteristic. B-type stars, which are non-supergiants on the main sequence, may become suddenly a Be star. The explanation of such a rapid transition is still unknown. Some professionals men- tioned the utility of such a survey. This survey can be conducted with even a spectrograph of R=1000. It should consist of recording spectra of B-type stars (subtypes IV or V) and would detect changes of Balmer emission lines.

6. Sharing the Data

The purpose of these amateur surveys is to provide to the scientific community regular and quality data to feed their star models and theories.

As we have seen for the first period of Be star monitoring, amateurs had acquired quite reliable data but this was not known by the professionals. When the data were published on the internet, some professionals found them and asked for our agreement to use them. So, putting data accessible to everyone on a web site is the first simple way to connect with the professional community. But it is not enough. A professional will not use data he does not trust. E-mails exchanges had to happen to describe the conditions of acquisition, and the processing done. Raw data were processed by the professional and then compared to calibrated spectrum published to check the quality of the data reduction made. Reliable software tools, algorithms and sequence of processings are very important to get credibility and give confidence to the professionals to use our data profile set.

Another way to get in contact with professional community is to actively contact them. We always received a nice welcome and got active support by the French Be star specialists.

In 2003, a specific seminar was organized by JP.Rozelot, a professional astronomer at the "Cote d'azur" observatory, Nice. Sponsored by the French Center of National Scientific Research (CNRS), amateurs and professionals gathered for 3 days of intensive communications around techniques and applications for spectrography.

In 2004, we decided to exploit the link between pros and amateurs for "science astronomy" by creating an informal international group ARAS (Astronomical Ring for Access to Spectrography) with an internet list and web site. Professionals are encouraged to join and make calls for observations. Internationally, amateurs who want to contribute can get the latest information for specific objects under a campaign of observation and even get explanations about the variations observed. We get direct satisfaction from our observations, by knowing it is contributing to science. We are now starting to organize specific programs, which are listed on the ARAS web site. Our next step will be to design a web database to provide free access to the spectrographic observations done by the ARAS members. This is NOT in direct competition with the AAVSO. Maybe one day
they will add a spectral database in addition to their photometric data.

However, there is still one question not answered: which data should be published?

Publishing the raw images is not appropriate. To obtain the real spectrum profile, many calibration steps are involved and specific instrumentation corrections are required. The most standard data set is a spectrum profile, calibrated in wavelength. Normalization to the continuum and removal of atmospheric lines is still under debate. The data set format would ideally be under fits format. But as fits is not as standard as it should be, we prefer the basic ASCII format, a text file with two columns, one for wavelength, one for intensity. As there is no header, additional information like observation time, object name, location and instrument information should be provided separately.

One step further in the data reduction can be envisioned. One can publish the Equivalent Width data or Doppler shift for a specific star, provided that the reduction method is well described and reliable. For example, Ernst Pollman in Germany publishes regularly H-alpha EW data.

7. Results

Starting in 1990, the compilation of the T60 data set and the personal observations of Christian Buil covers 280 Be Stars. Some of the stars have more than 30 observations during this time period.

Delta Sco is a spectacular example of a B-type star suddenly transitioning to Be phase. It was first observed in May 2001 as exhibiting a strong Hα line in emission. This star continues to vary and the latest observation shows clearly an outburst.

Nu Gem – recently an evolution in the peak intensity on the blue side of the Hα line triggered the attention of our ARAS list. Coralie Neiner, a professional astronomer, sent us explanations as to why the red peak was stronger in 1999. "The observations could indicate that there is a one-armed oscillation of the disk, i.e. the disk is inhomogeneous and we see a denser part rotate around the star with a period of 8-9 years. This star is thus an interesting target for a long-term monitoring."

Fig 12 - Evolution of Delta Sco between August 2004 and February 2005 - Hα region - C11 and LHIRES2 spectrograph (sampling 0.115 angstroms).

Fig 13 – Nu Gem observations in September 1999 – Hα region – 8-inch telescope – sampling 0.93 angstroms per pixel – the peak on the red side is the strongest one.

Fig 14 – Nu Gem observation in March 2005 – Hα region – 11-inch telescope – sampling of 0.115 angstroms per pixel, LHIRES2 – the peak on the blue side is seen this time as the strongest one.

Zeta Tau - If you want to be infected with the study of Be stars, take a spectrum of Zeta Tau. Then wait one or two months, and take another one. That way you will get the virus, as Zeta Tau has always shown interesting profile changes. From February 1990 up to now, 72 spectra were recorded. This star is suspected to be a binary system and the analysis of the periodic variation of line profile features is used to build the system model. Many observations over a long period of time are necessary to detect all the periods of the complex system. It is amazing to see how a line profile can vary and change shape.

**HD206773** – This 7th magnitude star exhibited a strong decrease of the Hα line after a spectacular outburst. In one month the relative intensity went down by a factor of three and after three months the Hα line was almost invisible.

Fig 16 - A spectacular outburst of the Be star HD206773 in 2001. The spectrograph resolution is 8500 and the telescope is an 8-inch. (Takasahi CN-212).

The list of results is far from exhaustive, just a few outstanding examples are shown. Besides the passion of collecting data which clearly demonstrated the activity of our colleagues, the best reward is the publication in a professional article. A list of some of them are provided in the reference section to demonstrate once again that quality amateur work in the field of spectrography is possible.

8. **Conclusion**

The Be Star survey is only one example of what the spectrography technique can open for amateurs. With the appearance of commercially available spectrographs, "home-made" spectrograph drawings available on the web and free software tools which has been proven as reliable tool for scientific application, amateurs can really consider spectrography as an exciting "not-so-complicated" technique to add to their astronomical toolbox.

The field of observations is absolutely not limited to Be stars. Regular spectroscopic monitoring of all types of variables are of interest. One can mentioned here the study made by Dale Mais on the Mira-type stars, or our campaign at the T60 in 2004 on hot super-giants stars.

Stellar variability may not be observed with a photometric only survey. A line profile variation may not drive an overall star luminosity change. On the other hand, a spectroscopic survey may benefit photometric measurements, since a spectrum covers a small range of wavelength.

The science of Astronomy is a very exciting field, mastering new techniques to provide trustable data which can contribute to the better understanding of the Universe around us is not a dream. With spectrography as a new tool, even more opportunities will arise.

9. **Acknowledgements**

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Implementation of a Fully Automated Fiber Spectrograph

Thomas G. Kaye
Spectrashift.com
Prospect Heights, IL 60070
tom@airgun.com

Michael Schwartz
Tenagra Observatories
HC2 Box 292, Nogales, AZ 85621
mbs@tenagraobservatories.com

Abstract
The last decade has seen a revolution in automated telescopes. Off-the-shelf hardware, software and standardization have allowed amateurs to excel in this area. Here we present the mechanical and software implementation of a fiber fed spectrograph and the methodology to achieve fully automated operation. The requirement that the imaging science camera coexist simultaneously with the spectrograph feed involved the design of a novel fiber interface and custom software. First light data is presented highlighting software selection between high and low resolution modes. © 2005 Society for Astronomical Science.

1. Introduction
The revolution in non-professional astronomy continues with automation software allowing for previously unheard of capabilities. While professional astronomers tend to redevelop software for each new telescope, the non-professional community is seeing tremendous progress in working with standard scripting interfaces between commercial software packages. In many cases the software is ahead of the telescope hardware since no standardized telescope mount exists at this time. Dramatic new developments are more likely on the hardware side than in software.

The Tenagra II is a 0.8 meter fully automated RC design. As with many high performance mounts, great care was taken in the design to achieve excellent tracking over a several minute exposure without autoguiding. In 2004 the Spectrashift team and Tenagra Observatories established a collaboration agreement to implement Spectrashift’s tracking system and two spectrographs on the telescope. The challenge was to hold the star centered on the spectrograph fiber for up to an hour exposure.

The requirements for this system had several goals:
- No interference with the science camera.
- Simultaneous software switching between imaging and spectroscopy.
- Minimum one hour spectrum exposures with sub arc second tracking.
- No rework of the mount or the existing tracking system.
- Run two spectrographs off the same interface under software control.
- Fit the interface assembly inside the existing optical tube.

2. Equipment
The interface for the system is shown in figure 1. The black box is an AO-7 tip/tilt mirror from SBIG. The CCD camera is an ST-7 and the extra mount is for the fiber end (not shown). The assembly mounts to the access hole in the space originally allotted for an autoguider. A mounting arm holds the pickoff mirror within 2 inches of the optical axis. This allows the science camera an unrestricted view and the star of interest is moved off axis to the pickoff mirror. The star image is folded 90 degrees by the pickoff mirror and comes to a focus in front of the AO-7. From there the expanding light cone is reflected off the AO-7 mirror and into the first collimating lens (lenses not shown). A collimated beam travels a short distance to a second lens, which focuses the star on the fiber mounted at the far end of the assembly. A thin piece of glass is mounted in between the 2nd lens and the fiber mount at 45 degrees to reflect 3% of the stars light to the tracking ccd in the ST-7. This allows the system to track directly on the star being analyzed.
Figure 1. The interface assembly for the Tenagra telescope with AO-7 and ST-7.

Figure 2. The fiber layout from the telescope to each spectrograph.

Figure 3. The low resolution fiber spectrograph.

3. Software

The key goal of this system was automated collection of spectra. The Tenagra telescope works off a scripted list of target objects for the night. The software needed to seamlessly switch between science camera exposures and spectra as required. Direct script control of the AO-7 and ST-7 was accomplished with Maxim-DL.

Since the guide chip is effectively off axis and has a very small field of view, the first step in acquiring a candidate star is to calculate the offset from the center of the optical axis to the guide chip. The scope is commanded to move to the offset position and a science camera exposure is taken and plate solved for position. This verifies that the proper target star is on the guide chip and if this test fails, the scope is software “bumped” into the proper position.

Once the field of the science camera is in the proper position, the guide chip is exposed to find the position of the star. The star must be moved to a particular position on the guide chip in order for the AO-7 to acquire and start tracking. The mount is again bumped to move the star into the proper fiber zone on the guide chip. There are six possible fibers to align the star on as shown in figure 4. Depending on which spectrograph and resolution is desired, the fiber is setup in the target list. Once the star is inside the 8x8 pixel tracking box, the AO-7 is commanded to start guiding and the tip/tilt mirror actuates. The AO-7 mirror actuation moves the star to a pre-defined pixel location, which represents the center of a given fiber. The guiding computer then releases control back to the scripting computer, which will start the spectrograph exposure.
Short term tracking is handled by the AO-7 but over the long term, drift accumulates. As this happens, the tilt position of the AO-7 increases farther from the home position until it eventually runs out of travel and loses the star. In order to compensate for this, the main computer poles the guide computer watching the mirrors angle position. When it comes close to the limit, another bump is initiated by the mount to bring the star back on center. The AO-7 keeps guiding during this move and which does not affect the exposure. After the exposure is completed, the guide computer commands the AO-7 to stop tracking and home the mirror for the next acquisition.

4. Summary

To date all the processes necessary to acquire and track a star under script control have been completed. Initial tests show that the low resolution spectrograph can produce a spectrum from an 8.5 magnitude star in 5 minutes at 75-1 signal to noise. Best focus and fiber position have not been determined yet as those scripts are under development in the debug stage. Next step is to build the containment building for both spectrographs. The high resolution echelle spectrograph is under development at the Spectrashift facility and is expected to come on line in the fall of '05.

To the author’s knowledge, this the second fully automated spectroscopic telescope in the world and the only one to mix imaging with spectroscopy. The Fairborn Observatory in AZ currently hosts a 2 meter robotic telescope / spectrograph without imaging capability. The Tenagra telescope’s first planned observing mission will be to spectroscopically monitor Titan for volcanic activity in order for Cassini to do follow-ups. This project again illustrates the gains made at the non-professional level in telescope automation.

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Submitted Abstracts
Lowell Observatory’s  
Discovery Channel Telescope

Edward Bowell  
Lowell Observatory  
1400 W. Mars Hill Road  
Flagstaff AZ  86001

Abstract

Lowell Observatory is developing a 4.2-m wide-field telescope having significant capabilities for solar system and broad spectrum astronomical research. The Discovery Channel Telescope (DCT) will be able to switch between 2-deg-FOV imaging using a prime-focus camera and 30-arcmin-FOV instruments at Ritchey-Chrétien focus. The DCT is to be located 65 km SSE of Flagstaff, Arizona, at 2360m altitude. The site produces year-round median image quality of 0.84 arcsec, with a first-quartile averaging 0.61 arcsec. DCT will feature active optics and alignment capability, and the prime focus camera will contain 40 2k x 4k CCDs. The maximum RC instrument payload of about 2300 kg will be suitable for large instruments or suites of co-mounted instruments. Preliminary mechanical and optical design has been completed and a concept design review took place in July 2004. The mirror has been cast and sagged, and provision is being made for eventual provision of a Nasmyth focus. The telescope is being developed in partnership with Discovery Communications, Inc., which will use it and the association with Lowell Observatory to devise educational programming about astronomy and technology. First light is expected in 2009. Lowell Observatory is seeking additional partnerships for DCT. © 2005 Society for Astronomical Science.
Some Suggestions for Writing up Asteroid Lightcurve Observations for Publication

Alan W. Harris
Space Science Institute
4603 Orange Knoll Ave.
La Canada, CA 91011-3364

Abstract

Over the years I have spent many hours re-constructing photometric observations from published works that would have been much easier if the authors had provided full and accurate details of the observations being reported. In this paper I will present the essential elements of reporting that render the photometric observations reported most useful and convenient for future analysis. The underlying principle to keep in mind is that the composite lightcurve you construct may not be right, or may not serve the purposes of some future research project, such that someone may need to recover your original observations as a time series rather than just the composite lightcurve presented. Other ancillary information should be provided to assist in using the data in combination with observations at other times, e.g. the aspect data (sky position, phase angle, light time correction if applied, etc.), and the color band and magnitude scale information. In a composite lightcurve figure, vertical (magnitude) offsets applied night to night should be indicated (in the figure or in a data table), days of observation should be plotted with different symbols and identified, and if coverage on a single night exceeds the rotation period such that the time series is "wrapped", it should be possible to determine the actual time of observation of each datum. In constructing a composite, it is essential that the period used to fold the data be the exact value stated so that the exact observation times can be reconstructed from the composite. Ideally, all data presented should also be made available in electronic form in a public archive (this will be a topic of considerable discussion at the meeting). Another essential feature is to reference and critically evaluate any previous lightcurves of an object that can be found in the literature. In addition to checking for consistency of results and possibly incorporating other observations in the analysis, the task of reconstructing observations from ancient publications will serve by example (good or bad) the importance of unambiguous presentation of results, and what is required to achieve that. At the risk of committing a blunder myself, I will illustrate my presentation with an example of an "ideal publication" of a lightcurve result, done in the style of the Minor Planet Bulletin. © 2005 Society for Astronomical Science.
Recent Arecibo and Goldstone Radar Imaging of Near-Earth Asteroids

Lance A. M. Benner
Jet Propulsion Laboratory

Abstract

This talk will describe near-Earth asteroids (NEAs) recently imaged by radar at Arecibo and Goldstone. In the past year, 25 NEAs have been detected by radar, 19 of them for the first time. Highlights include new images of 25143 Itokawa (1998 SF36), the target of Japan's Hayabusa mission, which is scheduled to rendezvous with the asteroid this summer; 11066 Sigurd, a 3.5-km-long contact binary that is shaped like a peanut; 1998 ST27, a binary with a distantly orbiting and rapidly spinning satellite; 3908 Nyx, topographically rugged, kilometer-sized object made of basalt; and 2004 MN4, which in April 2029 will make the closest approach to Earth (only 5.7 Earth radii) ever predicted. © 2005 Society for Astronomical Science.
ASTEROID LIGHTCURVE RESULTS FROM FLORIDA GULF COAST UNIVERSITY

Thomas Bennett, Michael Fauerbach and Manuel J. Mon

Evelyn L. Egan Observatory
Florida Gulf Coast University, 10501 FGCU Boulevard, South, Ft. Myers, FL 33965-6565

Introduction:
Photometric lightcurve measurements of asteroids provide a great opportunity for small telescopes 10.2m to 40cm to become involved with useful astronomical research. Many amateurs and small university telescopes collaborate with professional observatories in this endeavor. The list of potential targets, and the reasons for studying them, are many. Even for uninterested, non-calibrated observations, there are numerous opportunities, like:
- Asteroids that are targets for radar observations or spacecraft flybys
- Asteroids for which no known lightcurve exists
- Studies of potentially hazardous asteroids
- Asteroids that have repeated periods with high synchronicity

These asteroids can be studied at shorter timescales to provide detailed information about the target asteroids. A list of potential target asteroids can be found at the Collaborative Asteroid Lightcurve Link (CALL) webpage [1] and associated links.

Asteroids with a known rotation period, but with a large uncertainty, or those which are targets for shape modeling efforts, are especially good targets for undergraduate students—both astronomy majors and non-majors—in research activities. For asteroids with a rotational period ~5th, full lightcurve coverage can be obtained in only once or twice a night. This way, undergraduate students can be involved in the entire observational process, from target selection, actual data taking, to the data analysis, within the limited amount of time they can spend on these projects.

Asteroid Gaine,

The Evelyn L. Egan Observatory is located on the campus of Florida Gulf Coast University (FGCU) in Fort Myers, Florida. It is home to a 16-inch Ritchey-Chrétien telescope from Optical Guidance System (OGS). The telescope is mounted on a robotic Paracoumo (GT100ME) and is computer-controlled, via Software Bisque’s TheSky software, for photometric measurements. The telescope is equipped with an OAG IFW filter wheel with Band BVRI filters, a high quantum efficiency, back-illuminated Andesite APT camera and a 0.63-m focal reducer, bringing the system to F5.5. The field of view of a single exposure using this setup is roughly 30 arcminutes by 20 arcminutes. Camera control and subsequent image calibration was performed via MaxImDL by Diffraction Limited. Photometric data was obtained utilizing the computer program "IPO CamPus" from B&W Publishing.

Results:
The photometric observing program at FGCU was started in the summer of 2004. Utilizing funds from Florida Space Grant Consortium’s "Undergraduate Space Research Participation Scholarship" we were able to purchase the program and to provide a scholarship for one student. Initially, the photometric program was centered on the photometric monitoring of asteroid Gaine, which was an interesting candidate for a photometric monitoring project.

More recently, we have been conducting photometric monitoring of the asteroids 550 Capella, 456 Hestia, and 2887 Pallas, with the aim to measure their brightness variations and to determine their physical properties, such as their albedo, size, and shape. The photometric data was obtained using the Evelyn L. Egan Observatory, located at the FGCU campus.

Periods with Large Uncertainties:
1360 Gaine was discovered on December 13, 1999 and was observed during 4 nights (January 13, 2000, February 13, 2000, March 13, 2000, and April 13, 2000). It was observed using the 16-inch Ritchey-Chrétien telescope at the Evelyn L. Egan Observatory. The asteroid's period was determined using the method of least squares and found to be 5.25 days with an amplitude of 0.05 mag.

References:

Acknowledgments:
Data from the COMET survey, which is supported by the European Space Agency (ESA), is used in this work. The data are accessible at the following URL: http://comet.org/

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