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CCD photometry and the stellar content of the central region of IC 1805

David Bretz

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CCD Photometry
and the Stellar Content of
the Central Region of IC 1805

by
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B.A. University of Rochester
(1987)

A thesis submitted in partial fulfillment of the
requirements for the degree of Master of Science
in the Center for Imaging Science at
the Rochester Institute of Technology

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CCD Photometry and the Stellar Content of the Central Region of IC 1805

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Abstract

Images of the central region of the open cluster IC 1805 were made with a 2033 x
2044 Kodak CCD imager using broad-band filters corresponding closely to the standard
B, V, R, I, system of Johnson. Calibrated magnitudes were measured for 136 stars in the
cluster, and temperatures and masses were estimated for 115 of these. A mass function
was derived as defined by Scalo (1986) and is compared with an earlier determination for
this cluster. The mass function also is compared with that of a similar star-forming
region in Cygnus and with the initial mass function of field stars. A brief overview of
CCD design and operation is included, along with some specifics about the R.I.T.
camera, image reduction, and photometry procedures.
This work is dedicated to my mother and father

Robert Lawrence Bretz
and
Alma Linda Bretz

for their generosity and support
and for provoking me to question things.
I wish to gratefully acknowledge the assistance and inspiration of Dr. Zoran Ninkov.

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1 INTRODUCTION

It is generally understood that stars condense from contracting molecular clouds to form tight clusters which gradually disperse. Of interest is the way the material in these protostellar clouds is distributed with respect to stellar mass at birth -- i.e., what is the proportion of high-mass stars formed to low-mass stars formed? Knowledge of this "initial mass function" (IMF) is key to understanding both stellar evolution and galactic evolution.

Determinations of the "field-star" IMF within our own galaxy have been made from star counts in the solar neighborhood, but this is an average over many contributing clusters, and requires assumptions to be made regarding stellar lifetimes (cf. Scalo 1986). Clusters in young OB associations offer measurements of the IMF just after star formation, before members have had time to migrate a long distance, and before fast-evolving high-mass stars have died. The most massive stars in these clusters lose mass at a high rate causing large surrounding volumes of obscuring gas and dust to be cleared away relatively quickly, thus making members easy to locate. Therefore, an estimate of the mass of a young OB cluster will be a relatively complete inventory of stars formed in one location during the same epoch. Studies of individual cluster IMFs can help answer questions regarding the variability of the galactic IMF.

Theories of stellar evolution imply that the mass of a star primarily governs both its luminosity and temperature, and determine how these two quantities will change over the lifetime of a star. Hence, measurements of stellar luminosities and temperatures provide indirect measurements of mass. Stars are very similar to blackbodies and thus, their
temperatures can be inferred from broad-band color differences measurable with a light-sensitive device such as a CCD camera.

At the time of this study, the R.I.T. CCD camera contained an uncoated Kodak KAF-4200 silicon CCD which uses two-phase buried channel processing. This device was characterized by Ninkov et al. (1993). Each of the 2033(H) x 2044(V) pixels are 9.0 μm square and photosensitive to wavelengths between 0.4 μm to 1.1 μm. The dark current of the device is specified to be less than 10 pA/cm² (50 e⁻/pixel/sec) at 25°C. When the device was cooled to -35°C for this study, the average dark current was measured to be 0.03 e⁻/pix/sec with 2.5% of the pixels having a dark current greater than 0.085 e⁻/sec. The read noise was approximately 15 e⁻ rms. Charge-storage capacity was limited to approximately 82,000 e⁻ at unit gain (20 e⁻/Analog-to-Digital-Unit), and approximately 20,000 e⁻ at an amplifier gain of 4 (5 e⁻/ADU). The output was digitized to 12-bits (4096 gray levels). Accidental failure of the device in 1993 lead to its replacement with a similar KAF-4200 chip and an upgrade of the analog-to-digital converter to 14-bit output.

2 OVERVIEW OF CCDS

2.1 DESIGN AND ARCHITECTURE

Charge-Coupled Devices (CCDs) were developed in the late 1960s and early 1970s from the work of Willard Boyle and George Smith (1970), who received the patent for the charge-coupling principle in 1974. CCDs have been used for film digitizers, video and still cameras, and for scientific, military and industrial applications. More efficient and more linear in response than photographic plates, they are now common tools in the
to larger format devices with lower thermal noise characteristics, such as the Kodak KAF-4200, which are useful for performing photometry on large numbers of stars without the need for cryogenic temperatures. The history and current state of astronomical CCD designs is discussed in more detail by McLean (1989) and Janesick & Elliot (1992).

2.1.1 Conduction Bands in Silicon

CCDs rely on the conductive properties of silicon atoms in a lattice (cf. Dereniak & Crowe, 1984). Electrons in atomic orbits around atoms are allowed only discrete energy levels near the nucleus. In a crystalline solid, such as silicon, the periodic spacing of atoms the allowed energy states, or orbitals, to split into multiple levels, and bonds are formed by the outermost electrons. This new band of energy states is called the valence band. The proximity of the atoms also causes multiple splitting of unoccupied states of higher energy, forming a conduction band. Conduction-band orbitals do not contribute to the bond between atoms, so electrons within the conduction band are free to roam throughout the solid. In metals, the conduction and valence bands overlap, meaning that the lowest energy level of the conduction band is lower than the highest energy level of the valence band (Figure 1a). Electrons in a metal need no extra energy to move into the conduction band and migrate. In an insulating material, the valence band is filled while the conduction band is empty and separated from the valence band by a large energy gap (Figure 1b). In semi-conducting material, such as pure silicon, the valence band is filled, but the energy gap separating it from the conduction band is small, allowing electrons to
move into the conduction band if given a sufficient, but small amount of energy, by thermal interactions or otherwise (Figure 1c).

A valence-band electron in silicon can absorb a photon if the energy of the photon is larger than the energy gap. When this happens, the electron enters the conduction band and is called a photoelectron, which may be counted. The positively charged vacancy left by the promoted electron is called a hole. In the absence of an electric field, the photoelectron wanders randomly through the silicon substrate until it recombines with a hole. To preserve the detection of the photon, the electron-hole pair must remain separated. In a CCD, this is accomplished by forming an array of metal-insulator-semiconductor (MIS) or metal-oxide-silicon (MOS) capacitors by placing metal electrode gates above a layer of insulating material on top of the substrate. A small positive voltage on one of the gates induces a depletion of holes near the surface where wandering photoelectrons will be held against the oxide interface. A two-dimensional array of
these storage sites forms the basic structure of a CCD imaging device. Each storage site is a picture element (pixel) in the resulting image. For more information on MOS capacitors, see Nicollean and Brews (1982) and Kim (1979).

2.1.2 Charge Clocking and Readout

To access the information contained at each pixel, the charge must be transferred along columns of adjacent pixels one row at a time to a serial register, where each row is read out pixel by pixel. The charge is transferred by raising and lowering the voltages on the electrode gates with timed clocking voltages to allow the charge to spill from one pixel to the next in a controlled manner.

Several transfer architectures can be employed: two-phase, three phase, four-phase, and virtual (one)-phase (cf. McLean, 1989). In a three-phase device, each pixel is overlaid with a set of three metal electrodes. A schematic diagram is shown in Figure 2. While integrating an exposure, the center electrode is held at a more positive voltage than the other two, creating a potential well where photoelectrons are collected. Charge may be moved to a neighboring electrode by raising its voltage, thus causing charge to diffuse across the region beneath the two raised potentials. After a time, the voltage on the original integrating electrode is decreased (made less positive), thus pushing the remaining electrons to the new storage region. The three electrodes at each pixel are operated simultaneously across the entire CCD, moving charges along columns of pixels through repeated pulsing (clocking) of the electrode voltages. Channel stops, or barriers made by placing a dopant in the silicon, prevent charge from leaking across columns.
Figure 2. Charge coupling and the timing waveforms for a three-phase CCD.
In a three-phase device, charge can be moved in the opposite direction by reversing the clocking pulses. Bi-directional transfer is not always necessary and a unidirectional charge transfer can be achieved with fewer electrodes. In a two-phase device, each pixel is defined by two electrodes and charge is diffused into one half of the substrate beneath each electrode (Figure 3). The implanted charge creates a deeper potential well beneath it forming a one-way step for each electrode to control. By pulsing one of the two electrodes high and low, while holding the other at an intermediate level, the charge is shuffled in the one direction. The R.I.T. camera is a two-phase device. Because one of the two electrodes in the two-phase device is always held constant, it can be replaced by an implanted charge (or a virtual phase) and only a single electrode is needed to move charge (Janesick, et al., 1981). A four-phase device uses four electrodes to create four potential steps which allows full control of phase combinations and the direction of charge transfer.

2.1.3 Buried-channel Devices

A CCD can be of a very simple design -- a p-type doped silicon substrate covered by an insulating material (SiO₂) with overlaid metal electrodes to store and manipulate the charges at the surface through the insulating layer. This is termed a surface-channel CCD. Unfortunately, the surfaces of crystal lattices have irregularities and defects which trap charges and release them slowly, thus degrading charge transfer. It is better to manipulate charges below the surface where the crystal structure is more uniform.

By including a doped n-type silicon layer between the existing p-type layer and the insulating layer (Figure 4) a p-n junction and a depletion region is formed within the
Figure 3. Two-phase CCD. Selective doping creates stepped potentials.

Figure 4. A single storage site in the buried-channel CCD. The collection layer lies below the surface where the gate depletion and the depletion region of the p-n junction overlap.
lattice, attracting electrons to the depleted n-side and drawing holes to the negatively ionized p-side. The width of the p-n depletion region can be increased by applying a positive bias to the n-side (called back biasing) which partially overlaps the electrode-induced depletion region created at the surface. The most positive potential, called a buried channel, will then lie within the overlap region below the surface near the electrode gate, rather than at the surface. Almost all modern CCDs, including the KAF-4200, are designed this way, although it is more typical for the p-type and n-type layers to be reversed and the majority charge carriers to be the positive holes rather than the negative electrons (Walden et al., 1972). Buried-channel devices support higher clocking frequencies and have higher charge-transfer efficiencies, but cannot be operated at temperatures below 77-K (Dereniak & Crowe, 1984), and have lower charge-storage capacities (McLean, 1989).

2.2 SPECTRAL RESPONSE

The spectral response of a semiconducting material is determined by the energy gap ($E_g$) which separates the valence band from the conduction band. Only photons with energies equal to or greater than $E_g$ are capable of exciting electrons across the band gap, and so photons are absorbed only if their wavelength is less than the cutoff wavelength, $\lambda_c$:

$$\lambda_c = \frac{hc}{E_g}$$  \hspace{1cm} (1)

will be absorbed (Dereniak & Crowe, 1984). This cutoff wavelength is the limit of red sensitivity. Silicon has a band gap of $E_g = 1.12$ eV, so that $\lambda_c = 1.1$ μm. The spectral sensitivity of the KAF-4200 is shown in Figure 5.
Figure 5. A typical response curve for the KAF-4200 from the Eastman Kodak Performance Specification. Actual response varies somewhat between chips.
2.2.1 Front Illumination vs. Back Illumination

The high-frequency (short-wavelength) sensitivity of silicon CCDs is limited primarily by reflection from the surface and the short absorption depth of these photons in the active medium. This is affected greatly by whether the device is illuminated from the front or from the back side.

The front side of a CCD is defined as that upon which the network of electrodes lies. The KAF-4200 is illuminated from the front, so the incident light must pass through the material making up this network, which is typically thicker than the absorption depth of blue (400nm) photons. Thus, the majority of blue photons never reach the active substrate. An increase in blue response can be obtained by flipping the device over so that it is illuminated from the substrate side (Shortes, 1974). Because of the shallow absorption depth of blue light, the silicon must be thinned to 10-30 μm so that the back surface is close enough to the depletion region created by the electrodes on the opposite surface. However if the substrate is too thin, red photons will pass through without being absorbed, making the thickness of back-side illuminated devices critical.

2.2.2 UV Coatings

Another way to improve the blue sensitivity of CCDs and extend the spectral response is to add a phosphorescent coating, such as lumogen, to the surface of the device which will absorb UV photons and re-emit the energy at longer wavelengths (Kristianpoller & Dutton, 1964). Preliminary analysis of data collected with such a device (Deeg & Ninkov, 1994) indicates that the response can be nonlinear at low light levels, and varies across the chip.
2.3 CCD OPERATION

Before photometric measurements of stars can be made from CCD images, it is necessary to remove, artifact signals which do not represent photon counts. These include the bias level, accumulated thermal noise (dark signal), and differences in sensitivity among pixels. There are some variations in the way these artifacts can be removed, with varying degrees of precision in the end result (cf. Newberry, 1991). For example, when images are divided, the output images should be saved with more bits per pixel (e.g. 32-bit-real if the inputs were 16-bit-integer) to avoid a loss of precision caused by a re-quantization. Other reduction methods are dictated by the way observations have been conducted and will be limited by the number and quality of available calibration frames. The reader is cautioned that the methods outlined here are not the most rigorous, but are more than adequate for the goal of this study: to resolve the mass function of IC 1805. All of the image reduction described in this paper was performed using IRAF (version 2.10).

2.3.1 Bias Level

The photoelectrons generated within each pixel of a CCD are sequentially transferred to a capacitor to produce an analog output voltage. This voltage measurement is digitized in an analog-to-digital (A/D) converter. The charge-to-voltage conversion adds a quantization error called read noise, which is bipolar and uniformly distributed. Most analog-to-digital converters accept only positive voltages. Thus, to avoid clipping off the bottom end of signals that are lower than the width of the read noise distribution, a positive bias voltage, or bias level, several times width of the read noise is introduced
into every pixel before each exposure. This bias level must be removed from each CCD exposure as a primary step in analysis.

If the bias level introduced before each exposure were identical for all pixels on the chip, then subtraction of a single number from every pixel would remove the bias. Unfortunately, the bias level is not uniform across the device. However, it typically has a consistent structure which can include column banding (slight variations from column to column) or ramping (an increase toward the corner closest to the output amplifier). If this bias pattern is significant in relation to the variation caused by read noise, the bias must be measured by reading the array without integrating any charge and then subtracted from every observation.

A single measurement of the bias includes the read noise from the output amplifier. This noise is a random error in each pixel value that generally is independent of pixel position, and ideally has a gaussian distribution (with mean \( \mu = 0 \), and variance \( \sigma^2 \)). Every observation frame also contains this read noise. Subtraction of a single bias frame from an observation frame doubles the read-noise variance at each pixel, and thus increases the total noise by a factor of \( \sqrt{2} \). That is:

\[
\sigma_{\text{total}}^2 = \sigma_{\text{read}}^2 + \sigma_{\text{read}}^2 = 2\sigma_{\text{read}}^2
\]

and thus:

\[
\sigma_{\text{total}} = \sigma_{\text{read}} \sqrt{2} = 1.414\sigma_{\text{read}}
\]

To avoid substantial increases in the noise within an image when removing the bias, a master bias frame should be created and used for bias removal by computing the mean or median of several bias frames. This is desirable because the standard deviation of \( N \) measurements of a random variable with a gaussian distribution decreases by the factor
\(\sqrt{N}\) (Dougherty, 1990). As an example, if 16 bias frames each with noise \(\sigma_{\text{read}}\) are averaged, the variance of the mean bias frame is \(1/16\) as large as that for the individual frames. When the mean bias is subtracted from an object frame, the total variance due to read noise in the output will be:

\[
\sigma_{\text{total}}^2 = \sigma_{\text{read}}^2 + \frac{\sigma_{\text{read}}^2}{16} = 1.0625\sigma_{\text{read}}^2
\]

and thus the total noise will be:

\[
\sigma_{\text{total}} = \sqrt{1.0625\sigma_{\text{read}}^2} = \sigma_{\text{read}}\sqrt{1.0625} = 1.038\sigma_{\text{read}}
\]

The choice of average (mean or median) depends on the characteristics of the noise in the individual bias. An illustration of the reduction in rms error can be seen in Figure 6, which shows the histogram of the central 1000 x 1000 pixels from each of the following: a single bias frame; the median for each pixel of 7 bias frames; the mean for each pixel of the same 7 frames. However, a noise component that is non-stationary, non-stochastic, or with a non-zero mean value, such as is generated by cosmic ray events or electronic interference, will cause some pixels to have values far from the mean. When such noise is present, the median becomes the desired combining method because it is less corrupted by these outlying values (Frieden, 1983).

In addition to spatial variations, the bias level also can vary in time (bias drift). This means that bias frames should be measured throughout a night of observations. If the bias level variations in time are small, then the mean or median of all available bias frames gives the best results. If the bias level varies significantly over time, one can either use only those bias frames near enough in time to the observations to provide a stable set, or one can measure the drift and account for it. This involves calculating independently the bias structure (which is constant) and the bias level as a function of
Figure 6. Histograms of the central 1k x 1k region of: a single bias frame, the median of 7 frames, and the mean of 7 frames.
time. A full description of this method can be found in the user’s guide for the software package MIRA, developed by Michael Newberry and distributed by Axiom Research Inc. The method can also be used to account for drifts in dark signal.

2.3.2 Thermal Noise

If a CCD is allowed to integrate charge in the dark, some charge above the bias level will accumulate that is proportional to the temperature of the silicon chip and to the duration of the integration. These dark counts are the result of electrons that are thermally excited into the conduction band. McLean (1989) gives the formula for dark-current generation in silicon as:

$$N_D = \frac{I_0 d^2}{e^- \exp \left[ \frac{-(T_0 - T)}{9.96} \right]}$$

(4)

where:

- $I_0$ = the dark current in amps per square centimeter at room temperature ($T_0$)
- $d$ = the pixel pitch (the pixel separation, equivalent to pixel width for a 100% fill factor) in centimeters
- $T$ = the operating temperature in Kelvins
- $e^-$ = the charge on an electron ($1.6 \times 10^{19}$ Coulombs)

When taking long exposures, it is necessary to suppress the dark current as much as possible. Equation 4 indicates that the dark current drops by 37% for each 10-K drop in temperature. Typically, the chip is placed in a vacuum chamber and thermally connected to a reservoir of liquid nitrogen at 77-K = -196°C. The nitrogen reservoir must be replenished periodically, as it eventually evaporates away. The Kodak KAF-4200 chip has such a low dark current (<10pA/cm$^2$ at 25°C) that it can be successfully operated for
astronomical purposes in the -35°C to -60°C range. These temperatures are attainable with a thermoelectric Peltier cooling system, thus eliminating the need for cumbersome cryogenic liquids. The Photometrics CH-250 camera head on the R.I.T. camera utilizes a three-stage thermoelectric cooler and a mixture of ethyl-glycol and water which is circulated with an external pump and acts as a heat sink.

Dark current accumulates during all integrations and must be measured and removed if significant. It has the undesirable effect of reducing the well depth of each pixel, thus limiting the maximum possible signal-to-noise ratio. Also, it is necessary to hold the temperature of the CCD chip within a few degrees to minimize fluctuations in the dark current.

When the CCD is fabricated, unavoidable variations in the doping constituents lead to variations in the dark current among pixels. High-dark-current ("hot") pixels appear brighter than surrounding low-dark-current ("cool") pixels (Figure 7a). The histogram of an entire 600-second dark frame taken at -35°C is shown in Figure 8. Examination of a set of dark frames from the device revealed a cluster of three very hot pixels near one side which would saturate after only 100 seconds.

The dark current fluctuates due to its stochastic nature, and contains the ever-present effects of read noise. Thus, as with bias frames, it is desirable to combine several dark frames in order to improve the signal-to-noise ratio over that of a single measurement. When possible, the median is the preferred method for combining dark to eliminate isolated bright pixels caused by the absorption of cosmic rays. These are high-energy charged particles produced from the interaction of earth’s atmosphere with the solar wind and are absorbed at random times and locations, producing localized bursts of charge that
Figure 7. Two 100 x 100 pixel regions of a 600-second dark frame (T=−35°C) showing the horizontal smear caused by low charge-transfer efficiency in the readout register. Figure (a) is centered on column 150 and figure (b) is centered on column 1875.
Figure 8. Histogram of a 600-second dark frame with bias and cosmic rays removed. $T = -35^\circ C$. 
accumulate during long integrations. These charges are added to valid signal and may increase the accumulated charge to very high values. The isolated values are removed by the median filter.

2.3.3 Flat Field

Variations in the thickness of the substrate and the doping process during manufacture induce variations in sensitivity from pixel to pixel in the CCD. Shadows and other effects related to the optical system placed in front of the CCD may also cause sensitivity variations. These are corrected by dividing each object frame (after bias and dark subtraction) by an image of uniform illumination called a “flat field”. Flat-field images are generated by imaging a white surface inside the telescope dome (dome flats), by imaging the twilight sky (twilight flats), or by forming the median of several offset images of the night sky (sky flats). There are advantages and disadvantages to each method and some disagreement about which is best (cf. Gilliland, 1992; Gullixson, 1992; Massey & Jacoby, 1992). Imaging of the night sky consumes valuable observation time, but does provide the correct spectral illumination for dealing with sky-limited, faint-object work. The twilight sky provides a bright source, but is inconveniently brief and contains strong atmospheric emission lines. Dome flats are convenient, but a considerable effort must be made beforehand to ensure that the lights and reflective surface provide illumination that is sufficiently uniform both spatially and spectrally. Additionally, the detector and optics must be well shielded from stray light leaks, especially for twilight and dome flats.
Figure 9. A grayscale image and histogram showing the variation in the central 1k x 1k region of a B filter flat field divided by a Z filter flat field.
Because pixel response is wavelength dependent to some degree, a flat field image must be taken for each filter used during observations. A demonstration of the wavelength dependence was conducted using images of a reflective barium sulfate screen that was illuminated by daylight-corrected light bulbs. The screen was imaged with different broadband filters placed in front of the CCD, one blue (2mm of GG385 + 1mm of BG12 + 2mm of BG39 with a center wavelength of 430 nm and a FWHM of 80 nm, called a B filter) and one infrared (3 mm of RG830 Schott glass and 2 mm of WG305 with a center wavelength above 850nm, which we call a Z filter) while the CCD was mounted at the Cassegrain focus of the 24-inch telescope at the Mees observatory. After subtracting the mean bias from each, the ratio of the two images was computed to highlight spatial variations in the response difference. The histogram of the central 1k x 1k region of the ratio image is shown in Figure 9, along with an image of this region. The majority of pixels cluster around a single value that is the ratio of the average count levels of the two frames. The width of the distribution of this ratio is about 7% over the $10^6$ pixels, which is much larger than the variations caused by read noise.

2.4 ADVANTAGES OF CCDS

The greatest advantages of using CCDs for astronomical imaging are their ability to convert incident photons into detected events with high efficiency, and the linearity of response over a wide dynamic range. The KAF-4200 has a quantum efficiency between 20 and 45 percent for $\lambda = 500 - 900$ nm wavelength range (Figure 5). This is more than ten times larger than the quantum efficiency of most photographic emulsions (McLean, 1989).
A linear response implies that the measured number of electrons is directly proportional to the number of incident photons. Thus, when comparing similar exposures, an object that is half as bright as another will produce half the number of output counts. Equally, an exposure of an object that is half the duration of another exposure of the same object will again produce half the number of output counts. If a device’s response is not linear, the magnitudes of dim stars cannot be extrapolated directly from the magnitudes of bright standard stars. In such a case, the CCD response as a function of incident light would need to be accurately measured and corrected.

Another advantage of using a CCD is that the output is a digital electronic signal. This allows direct storage and analysis in digital computers, providing the user with an array of powerful tools for measuring distances, comparing fluxes, and performing complex image enhancements such as smoothing, sharpening, and deconvolution.

2.5 DISADVANTAGES OF CCDS

2.5.1 Bad Pixels

A single pixel in the CCD array with a defect will permanently affect the output of all the pixels whose charge is transferred through it. Thus, a non-functional pixel will block all charge in the column above that pixel, leaving a dark streak in the image. If a pixel in the output register is damaged, most or all of the array could become useless.

2.5.2 Blooming along Columns

Another undesirable effect found in CCDs is called charge blooming which is the leaking of charge from a saturated pixel to bordering pixels. Channel stops prevent
charge from spilling across rows, but since charges must be allowed to move along columns, similar barriers cannot be placed between these pixels. Thus, when a pixel is oversaturated, charge may spill over to the adjacent pixels in a column and corrupt their output.

An alternative to the CCD architecture is the charge injection device (CID) which overcomes some of these problems (Zarnowski, et al. 1993). In a CID, the charge at a pixel is read out independently of all others, removing the need to transfer all charges through other storage sites to one location. If a pixel becomes damaged by interaction with a cosmic ray, other pixels are not affected, giving CIDs an advantage for use in space-borne applications. Blooming does not occur because channel stops completely surround a CID pixel and over-saturation charges are drained off. In addition, the size of an array is not limited by the efficiency of charge transfer.

2.5.3 Limitations on Chip Size

Currently, the size of a CCD array is limited by the size of defect free silicon wafers, and by the need for quicker (or more efficient) readout schemes and extremely high charge-transfer efficiencies (CTEs). In a 2033 x 2044 CCD, the charge from the pixel farthest from the readout must undergo 4077 transfers. To ensure that a significant amount of charge is not left behind in the intermediate pixels, the transfer efficiency must be better than 0.99995 (Gilliland, 1992). Even at this efficiency, only $0.99995^{4077} = 0.82$ of the charge in this last pixel reaches the readout amplifier; the remainder is left smeared behind. Recent attempts to overcome this size constraint have focused on developing mosaics of four or more 2k x 2k devices (Janesick & Elliot, 1992).
2.5.4 Decreased CTE with Low Temperatures

Charge-transfer efficiency depends on the ability of thermal energy to move charges toward the lowest potential during clocking, so when a CCD is cooled to suppress the dark current, the CTE will suffer. The result is a smearing of the image in the direction opposite to charge transfer, which increases with distance from the readout. This effect has been observed in dark integrations which show point-like spikes due to "hot" pixels having above-average thermal noise. Figure 7a shows a 100 x 100 pixel region of a dark frame near the readout amplifier at a chip temperature of \( T = -35 ^\circ \text{C} \), while Figure 7b shows a similar region of the same dark frame, but on the other side of the chip far from the readout. In the latter, the large counts represented by bright pixels are elongated with "tails" along the horizontal rows. The effect is more noticeable at lower operating temperatures. The horizontal nature of the smearing indicates that it is occurring in the horizontal readout register at the "bottom" of the chip, where an entire row of pixels is transferred onto the output amplifier for each single pixel transfer of the columns.

2.5.5 Charge Trapping and Residual Images

It was noted in section 2.1.3 that buried-channel CCDs have far fewer charge traps than surface channel devices. However, there are indications that some trapping occurs in the KAF-4200 (Ninkov et al. 1993). The amount of trapped charge is believed to be very small and has only been detected where the pixels are severely oversaturated. The effect was found upon examination of a 300-second dark frame which was immediately preceded by a star-field exposure whose brightest component was oversaturated by a
Figure 10. Illustration of charge trapping. Figure (a) is a 300-second image of a star field taken through the Z filter. Figure (b) is a 300-second dark exposure taken immediately after, showing residual images of the saturated stars seen in (a). The images are the result of a 4 x 4 block average of 484 x 524 physical pixels.
factor of approximately five. **Figure 10a** shows a detail of the saturated star field and **Figure 10b** shows the residual image. The operating temperature of the chip at this time was -35°C and the gain was 4 (5e^-/ADU). The measured fluxes of the residuals are approximately 0.1% of the original image fluxes. The fact that a bias frame taken between these exposures showed no residual image indicates that the traps were releasing charge slowly.

The initial exposure was made through an infrared filter (the "Z" filter, section 2.3.3) used to restrict the incoming light primarily to wavelengths longer than 850 nm. A similar set of exposures involving a blue (B) filter (section 2.3.3) showed no clear residual. This would be expected, considering that blue photons are absorbed in a thin region nearer the surface of the silicon, whereas the red photons penetrate to a deeper volume where more defects are likely to trap the carriers.

A test of the phenomenon was carried out for both B and "Z" filters at an operating temperature of -60°C and no residuals were seen. Tests of the chip at room temperature made by the Eastman Kodak Company indicate no residual images were found. Therefore, it would seem that at very low temperatures the trapped carriers do not have enough thermal energy to escape from the traps, while at room temperature the carriers have enough thermal energy to escape easily and quickly.

3 **IC 1805**

3.1 **BACKGROUND**

IC 1805 (2h32m42s.36, +61°27m21s.6, 2000) is the central cluster of the Cas OB6 association and is contained within the HII region W4 that is located in the Perseus arm
of the Galaxy (at galactic longitude $l = 134^\circ.7$, and galactic latitude $b = +0^\circ.9$). The cluster is very young and has been classified as Trumpler class II3mN (Lynga, 1984). The core of the cluster contains the O5IIIIf star HD 15558 (SAO 12326) that is contained in the multiple system ADS 1920. Stellar radiation from the central stars has lead to a clearing of the central region as evidenced in the 4170 MHz maps of Akabane et al. (1967), that show a shell-like filamentary structure surrounding the core.

Previous photographic studies of IC 1805 have been carried out by Hoag et al. (1961), Ishida (1969), Moffat & Vogt (1974), and others. Spectroscopic studies include those by Hoag & Applequist (1965), Ishida (1970), and Walborn (1972, 1973). Vasilevskis, Sanders & van Altena (1965, VSA henceforth) published a membership study based on an analysis of proper motions which was revised by Sanders (1972). Recent and more accurate photoelectric studies have been conducted by Joshi & Sagar (1983, J&S henceforth), Sagar et al. (1986), Guetter & Vrba (1989, G&V henceforth) and Sagar & Qian (1990). Both J&S and G&V measured variable extinction in B-V across the cluster ranging from 0.68 to 1.4 magnitudes, however, the range is much smaller for the central 7 arcminutes (0.7 to 0.9) with a mean $E_{(B-V)}$ of 0.8 magnitudes. Both found a distance modulus of 11.9 magnitudes for the cluster (corresponding to a distance of 2.4 kpc) and determined an age of approximately $10^6$ years. Star counts made by G&V indicate an angular diameter for the cluster of 40 arcminutes or more, yet all of the O stars and half of the B stars are found within twenty arcminutes of the centrally located star VSA 165. G&V also found the ratio of total to selective absorption to be normal ($R_V = 3.1 \pm 0.1$) in the direction of the cluster, a result supported by the near-infrared study of Sagar & Qian (1990).
3.2 OBSERVATIONS

Images of IC 1805 were obtained using the R.I.T. CCD camera mounted on the Cassegrain 24-inch Northern Planetary Patrol Telescope located at Mauna Kea Summit, Hawaii on the night of 3 November, 1991. The observing team consisted of Zoran Ninkov and Roger L. Easton of the Center for Imaging Science at R.I.T., William J. Forrest of the University of Rochester, and Mark Shure of the University of Hawaii. The temperature of the chip was regulated to operate at -35°C, and the gain on the chip was set so that one AD unit (1 output count) was equivalent to 5 electrons of charge on the output of the CCD. The read noise of the on-chip amplifier and readout electronics combined was approximately 15 electrons. By using positions for stars in the IC 1805 field that have positional information given in the Hubble Guide Star Catalog, the scale was found to be 0.224 arcseconds per pixel. This made the field of view 7.5 arcminutes on a side. Small shifts in pointing between filters, and the exclusion of stars within 50 pixels of the edge reduced the effective field of view to 7.0 arcminutes on each side.

Observations centered on HD 15558 (SAO 12326, ADS 1920) were made through Bessell B, V, R, I filters in the Cousins system (Bessel, 1979). The observations were calibrated to the Johnson system on the assumption that the differences that exist between the two systems in the R and I passbands would lead to acceptably small errors. A synopsis of the observations is presented in Table 1. With each filter, a series of exposures of increasing duration was taken to observe a wider dynamic range than was possible from a single exposure. Figure 11 shows the 300-second R-band image of IC 1805 with the target stars numbered. Figure 12 is a detail from the same image which
Figure 11. 300-second R-band image of IC1805 showing target stars. The contrast has been reversed, so bright stars appear black.
Figure 12. Detail of 300-second R-band image showing central O5 star (#80) and possible companions.
Table 1. Observations of IC 1805

<table>
<thead>
<tr>
<th></th>
<th>B</th>
<th>V</th>
<th>R</th>
<th>I</th>
</tr>
</thead>
<tbody>
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</tr>
</tbody>
</table>

shows the central O5 star ADS 1920 and three possible companions more clearly. In addition to the observations, bias frames, dark frames, and flat fields were collected.

3.3 FRAME REDUCTIONS

3.3.1 Bias Removal

Only three bias frames were created on the night IC1805 was observed. Though the bias drift cannot be determined accurately, the bias level appeared to be sufficiently stable at a value near 25 ADUs. Because no irregularities, such as those caused by cosmic rays or electrical interference, were found in these three frames, a mean bias frame was created as a master and subtracted from each object frame.

3.3.2 Thermal Noise Removal

3.3.2.1 Dark Measures

A total of six dark frames were taken on the night of the observations. One was at a different amplifier gain setting, and could not be used. Of the others, two were 300-second integrations and three were 600-second integrations. The temperature of the chip was held at $-35^\circ$C with the aid of a thermostat contained in the thermoelectric cooling system developed by Photometrics.
The 300-second darks were intended to be used for reducing the IC 1805 data, as they were of the same duration as the longest of the object frames and had been taken immediately following the observations. However, the discovery of residual images in the first of these excluded its use, leaving only one useful 300-second dark frame. Although the three 600-second dark frames were taken several hours later, the temperature of the chip was stable enough to make these frames appropriate as dark measures. A 600-second dark frame has a greater signal-to-noise ratio than a 300-second dark frame. Also, the combination of three dark frames yields a better signal-to-noise ratio than one frame, and helps in removing cosmic ray events. No residual image effects were found in any of the 600-second frames. The mean of the 600-second frames was computed after removal of the mean bias and cosmic rays (described below). This mean dark frame was scaled to the object exposure time and subtracted from the 100-second and 300-second object frames. Because the dark current accumulated on the shorter object exposures was insignificant, no dark removal was performed on them.

3.3.2.2 Cosmic Ray Removal from Dark Frames

As stated in section 2.3.2, cosmic rays periodically strike the CCD and result in point-like images that accumulate during all integrations, even when the shutter is closed during dark integrations. The easiest way to remove cosmic ray events from darks is to form the median of three or more frames. Because initial data reduction used two 300-second dark frames, a method was developed for deleting the cosmic ray events and allowing an average of the two darks to be formed. This method computed the difference of the two darks (to remove the stationary dark signal) and preserved those features that
exceeded the read noise by five times the standard deviation of the noise. The \textit{imreplace} function in IRAF was used. The result is an image of the cosmic rays from each frame which can then be subtracted from its original. The success of this method lead to its use for removing cosmic rays from the three 600-second dark frames as well, even though a median filter could have been used.

3.3.3 \textit{Flat fields}

Sky flats in each filter were taken on the morning prior to the observations with the telescope at the zenith. The bias was removed from the flat-field images by using the average of two bias frames taken at the time the flats were taken. Far from being uniform and blank, the flat fields exhibited a number of features in each bandpass (Figure 13). There are a few very small (diameter $< 10$ pixel) very dark spots which were most likely caused by dust lying directly on the chip. No star images fell close enough to any of these to be affected. There are numerous large (diameter $= 100$-$200$ pixels) ring-like features of 10\% lower intensity than the surrounding level. These are attributed to shadows caused by dust on the vacuum side of a quartz window ahead of the CCD. Also, the western edge of the CCD frame was vignetted during observations that night. All of these effects produced localized reductions in the signal-to-noise ratio. Only a handful of stars (noted in Table 4) were significantly affected by the vignetted area.

3.3.4 \textit{Cosmic Ray Removal from Object Frames}

Section 3.3.2.2 contains a description of how cosmic rays can be removed from dark frames when multiple dark frames are available. Removal of cosmic rays from object
Figure 13. V-band flat field of twilight sky.
frames using these methods is possible only when multiple images of equal duration are available and can be properly aligned. No two cluster images in this study were of equal exposure, so a cosmic-ray detection and elimination filter routine (cosmicray in IRAF) was applied to each object frame after flat-fielding.

The routine cosmicray reduces the large peak values caused by cosmic ray events, but is unable to remove completely all associated charges. It operates by computing the ratio of the pixel value and the average of the surrounding pixels (excluding the brightest value) in either a 5 x 5 or 7 x 7 window. When this ratio exceeds a threshold set by the user, the pixel value is replaced with the average of the surrounding four pixels. The threshold value is important, if set too low, the routine will detect peaks in the base-level noise. This routine was used one or more times on each image with both window sizes before performing photometry. The filter should have been able to detect cosmic-ray events that overlapped with star profiles, but the number of such events is uncertain. For events that did not overlap with star profiles, the routine was only able to reduce their peaks to values just under the read noise. Thus, some visible residuals always remained which were necessarily avoided during photometry. The greatest benefit from removing the cosmic-ray peaks is that star-finding routines will produce fewer false detections.

3.4 PHOTOMETRY

3.4.1 Digital Aperture Photometry

In aperture photometry, the light from a single star is measured by isolating that star with a small aperture during observation. The method used here is called, synthetic
**digital aperture photometry**, where an aperture is projected onto the digital star image and the counts are summed from every pixel within that aperture.

The average background sky brightness within each CCD frame was estimated and subtracted by using 9-16 circular areas (depending on the frame) 40 pixels in radius that were void of stars. Local corrections to this mean sky estimate were later applied individually to stars using a sky annulus (typically 10 pixels wide) located just outside the point spread function of each star.

Star centers were located in each frame by using the IRAF routine *daofind*. This task fits a gaussian profile to possible stars with some user-defined parameters. The full-width-at-half-maximum (FWHM) of the point-spread-function (PSF) was approximately 10 pixels (2.2 arcsec) with notable elongation in the east/west direction in the longer exposures due to error in telescope tracking.

Variations in the spectral sensitivity of the pixels, resulted in a difference in the number of stars that could be detected in the longest exposures in each band pass. The 300-second R frame, which had very good signal-to-noise ratio, was taken to represent the field by numbering the 138 stars found in it. Tracking errors caused some of the 138 stars near the edge of this frame to be absent in the longest B, V, and I exposures. The low sensitivity of the chip to blue photons coupled with the reddening of all stars to meant that B magnitudes could be measured for only 51 stars.

Photometry was performed using the *apphot* routine in IRAF. Eleven apertures were drawn around each star with radii ranging from 10 to 60 pixels. This range allowed an appropriate aperture and sky annulus to be selected for each star. Preliminary tests showed that an aperture of 20 pixels in radius was sufficient to contain the entire image of a very faint star before it blended with the noisy background; for the brightest stars, an
aperture of 40 pixels in radius was needed. Approximately one third of the stars could not accommodate a sky annulus due to crowding, and a few bad cases required undersized apertures to avoid including counts from nearby stars. One dim star (#78) was so badly overlapped by the central star that a satisfactory photometry could not be obtained in any color.

3.4.2 Calibration of Instrument Counts

The count within an aperture centered on a star is a measure of the brightness of that star (assuming the measured counts include only those generated by incident photons). To be of use, this instrument count must be transformed onto a common scale known as the magnitude scale. The magnitude scale is a logarithmic scale related to brightness in the following way:

\[ m_1 - m_2 = 2.5 \log \left( \frac{b_2}{b_1} \right) \]  

where \( m_1 \) and \( m_2 \) represent the magnitudes, and \( b_1 \) and \( b_2 \) represent the brightnesses, respectively, of two stars. Because instrument counts \( c_1 \) and \( c_2 \) measured on a single exposure of these two stars are directly proportional to brightness (assuming the collecting device has a linear response), we can write:

\[ m_1 - m_2 = 2.5 \log \left( \frac{c_2}{c_1} \right) \]  

If an accepted value for \( m_1 \) is substituted, \( m_2 \) can be derived. If we substitute for \( m_1 \) from the literature and \( c_1 \) from the frame we wish to calibrate, but set \( m_2 = 0 \), \( c_2 \) can be derived as the count for a star of magnitude zero, or count-for-zero-magnitude (cfzm or czm):

\[ czm = c_1 10^{\left( \frac{m_1}{2.5} \right)} \]
Substituting the czm back into equation 5b we have a formula for the magnitude of any other star on the same exposure with the calibration star:

\[ m = 2.5 \log \left( \frac{\text{czm}}{c} \right) \]  \hspace{1cm} (5d)

If there are several calibration stars in a frame, an average czm can be calculated to reduce the uncertainty of the calibration.

If no calibration stars appear in the frame, it is customary to observe a few calibration stars in a nearby region. The czm is normalized to exposure time, if necessary, and becomes a count-per-second for a zero-magnitude-star. When the observations of IC 1805 were made, it was assumed that reliable photoelectric magnitudes in B, V, R, I would be available in the literature for the central star HD 15558 (SAO 12326) as well as for a few other bright stars in the field. For this reason, no standard stars outside the field were observed for calibration.

Photoelectric B and V magnitudes, and \( E_{(B-V)} \) values are quoted for ten stars in the field (not including the central star) by J&S (1983). G&V (1989) measured I magnitudes for eight of these stars and list spectral types for each, compiled from a variety of sources (Hoag & Applequist 1965, Ishida 1970, Walborn 1972,1973). The I magnitudes were measured in the Kron-Cousins filter system; they were transformed to the Johnson system as prescribed by Bessel (1979).

Unfortunately, however, no R magnitudes could be found in the literature for any star in the field, including the central star. To circumvent this problem, R magnitudes with the proper interstellar reddening were derived based on intrinsic color relationships for the nine stars of known spectral type, using the following standard formula:

\[ R = V - (V-R)_0 + (R_r - R_V)E_{(B-V)} \]  \hspace{1cm} (6)
(V-R)o is the intrinsic color index, unique to each spectral type, which was taken from Johnson (1966). V and E(B-V) were taken from J&S (1983). Sagar & Qian (1990) found that the interstellar extinction law toward the cluster is normal, so the ratios of total to selective absorption, Rr and Rv, were taken from Schmidt-Kaler (1982) to be 2.32 and 3.10, respectively.

Table 2. Calibration Stars

<table>
<thead>
<tr>
<th>Id. #</th>
<th>VSA#</th>
<th>BD #</th>
<th>SAO #</th>
<th>Sp. type</th>
<th>E(B-V)c</th>
<th>Bc</th>
<th>Vc</th>
<th>Rc</th>
<th>(V-R)c</th>
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<td>--</td>
<td>B3 Va</td>
<td>0.70</td>
<td>11.72</td>
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<td>10.77</td>
<td>0.685</td>
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<td>--</td>
<td>B2 Va</td>
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<td>0.755</td>
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<td>12324</td>
<td>O7 Vb</td>
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<td>9.58*</td>
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<td>0.675</td>
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<td>63</td>
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<td>B3 Va</td>
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<td>80</td>
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<td>O5 III(f)c</td>
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<tr>
<td>84</td>
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<td>B1 IVd</td>
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<td>11.15</td>
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<td>0.790</td>
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<td>A1 IV-Vg</td>
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<td>10.07</td>
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</table>

* These magnitudes were used for calculation of Mbol and T_eff but were not used for frame calibrations.
a) Ishida (1970)
b) Walborn (1973)
c) Walborn (1972)
d) Hoag & Applequist (1965)
e) Joshi & Sagar (1983) except for #80 (VSA 148) Vasilevskis et al. (1965)
f) Calculated using spectral type as described in text.
g) Guetter & Vrba (1989) magnitudes in the Kron-Cousins system.
h) Converted to Johnson system (cf. Bessel 1979)

Table 2 lists the stars used for calibration and the adopted values for their magnitudes. Problems with the photometry of certain stars (crowding, saturation, etc.), and the loss of others due to the small changes in the field of view, meant some of the published magnitudes could not be used for calibration, but were still needed to
determine the effective temperature and bolometric magnitude of a star. These magnitudes are noted in the table.

Using Table 2, calibrations of the 300-second B and V frames, the 100-second R frame, and the 20-second I frame were possible, eliminating the need for any of the shorter exposures. As stated above, a count-for-zero-magnitude (czm) was extrapolated for each calibration star in a frame and the average was used to calculate the magnitudes of all other stars in that frame. A czm was disregarded if it deviated more than $2\sigma$ from the average value for the frame. To calibrate the 300-second R and I frames, in which the original calibration stars were saturated, magnitudes for selected intermediately bright stars determined from the prior calibration were used. Table 3 lists the number of stars used to calibrate each frame and the estimated error (in magnitudes) associated with each calibration, based on the standard deviations of the czms.

Measured magnitudes for all stars in this study are listed in Table 4. Column 1 is a sequential identification number used in this paper, while columns 2 and 3 list the

<table>
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<th>Filter</th>
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<th>No. of Calibration Stars</th>
<th>RMS Error in Calibration (mags)</th>
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<td>I</td>
<td>300 sec</td>
<td>10*</td>
<td>.03</td>
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* The magnitudes for these stars were determined from the preceding (shorter exposure) calibration and the total rms error would include error from both calibrations.
Table 4. IC 1805 Magnitudes

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<th>Id. #</th>
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<th>VSA #</th>
<th>B</th>
<th>V</th>
<th>R</th>
<th>I</th>
<th>(M_{\text{bol}})</th>
<th>(\log(T_{\text{eff}}))</th>
<th>Sanders Prob. (%)</th>
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<td>--</td>
<td>--</td>
<td>blend with 77</td>
</tr>
<tr>
<td>75</td>
<td>13</td>
<td>185</td>
<td>12.16</td>
<td>11.59</td>
<td>11.04</td>
<td>10.56*</td>
<td>-1.78</td>
<td>4.28</td>
<td>43</td>
<td>near edge</td>
</tr>
<tr>
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<td>-</td>
<td>-</td>
<td>16.97</td>
<td>15.99</td>
<td>15.23</td>
<td>2.29</td>
<td>3.86</td>
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</tr>
<tr>
<td>77</td>
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<td>166</td>
<td>12.54</td>
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<td>11.50</td>
<td>11.07</td>
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<td>4.40</td>
<td>82</td>
<td>blend with 79</td>
</tr>
<tr>
<td>78</td>
<td>-</td>
<td>-</td>
<td>18.25</td>
<td>17.21</td>
<td>15.99</td>
<td>3.71</td>
<td>3.74</td>
<td>--</td>
<td>--</td>
<td>blend with 80</td>
</tr>
<tr>
<td>79</td>
<td>-</td>
<td>-</td>
<td>16.57</td>
<td>15.64</td>
<td>14.92</td>
<td>1.88</td>
<td>3.88</td>
<td>--</td>
<td>--</td>
<td>blend with 77</td>
</tr>
<tr>
<td>80</td>
<td>1</td>
<td>148</td>
<td>8.32*</td>
<td>7.83*</td>
<td>7.33*</td>
<td>6.87*</td>
<td>-9.19</td>
<td>4.40</td>
<td>2</td>
<td>saturated; blend with others</td>
</tr>
<tr>
<td>81</td>
<td>-</td>
<td>153</td>
<td>11.17</td>
<td>10.73</td>
<td>10.19</td>
<td>9.75</td>
<td>-6.53</td>
<td>4.44</td>
<td>2</td>
<td>blend with 80; listed as B2 by Ishida (1970); but photoelectric magnitudes were not available</td>
</tr>
</tbody>
</table>

(continued)
identification numbers assigned to some of the stars by VSA (1965) and Hoag et al. (1961) respectively. Columns 4-7 contain the derived B, V, R, and I magnitudes determined by this work. A colon following the tabulation in these columns indicates an associated photometric error larger than that quoted in the next section. Columns 8 and 9 list the bolometric magnitudes (Mbol) and effective temperatures (Teff) determined using the procedure outlined later in this paper. Column 10 lists the probability of the stars membership in this cluster based on the analysis of proper motions made by Sanders.
(1972). Column 11 lists comments that identify any special circumstances that may have affected the photometric measurement of that star.

3.4.3 Estimates of Photometric Uncertainties

The primary sources for uncertainties in the photometry were: (1) the calibration uncertainties described above which dominated for bright stars, (2) read noise, processing noise, and uncertainties in the estimated sky flux which dominated for the dimmer stars. Poisson noise due to the statistics in photon arrival ($\sigma = \sqrt{N}$) were found to be no more than 10% as large as these other sources. Uncertainty estimates based on the background noise (sky noise + read noise + processing noise) in each frame were calculated by measuring the variation in the sky counts within various apertures. To be precise, four groups of four closely spaced locations were chosen in starless areas on each processed frame. These coordinates were the centers of concentric apertures of 20, 30, and 40 pixels in radius. For each frame, the statistical variations in the number of counts within same-size apertures were tabulated for each group of four coordinates and for all sixteen. The deviations within each group of four overlapping apertures represented the uncertainty in photometry when a sky annulus correction was applied, while deviations among the four widely spaced groups represented the uncertainty when no annulus correction was applied. The latter was typically two to three times the former. These deviations, in counts, were added to model star counts covering the dynamic range of observed stars and converted to magnitude deviations. Figure 14 shows the estimated errors for the deepest frames in each bandpass based on sky annulus correction.
Figure 14. Estimated photometric uncertainties.
As expected, B measurements contained the greatest noise at a given magnitude. However, few stars dimmer than 16th magnitude in B were measured. The majority of measures were in V, R, and I passbands. Of these, the V magnitudes showed the lowest error for a given magnitude, but because of interstellar reddening and a lower V-band sensitivity (Figure 5), the signal-to-noise ratio was lowest in V for any given star. Therefore, a magnitude limit of 17.6 in V was imposed as a condition for “good” photometry.

3.5 CLUSTER MEMBERSHIP

In the absence of spectroscopic or kinematic information, the determination of cluster membership must be based on analysis of color relationships alone. Customarily, this is done, by plotting U-B vs. B-V for each star. In such a color-color diagram stars will be displaced from an intrinsic main-sequence position along a straight line of length proportional to the amount of reddening, distinguishing foreground stars of late spectral type and low reddening from cluster stars of early spectral type and high reddening. This works well only if the cluster members under study are earlier than spectral type A0. The stars in this study were expected to be as late as F and G.

The B-V vs. R-I and V-R vs. R-I plots, were used in an attempt to separate cluster members from field stars. The B-V vs. R-I plot is shown in Figure 15, and includes data for 49 stars for which B, V, R, and I magnitudes were available. The solid curve is the unreddened main-sequence taken from Johnson (1966) while the dotted curve has been reddened by the average extinction for the cluster, corresponding to an $E_{(B-V)} = 0.8$. Star #101, which is known to be a foreground A1 star with $E_{(B-V)} = 0.35$ (G&V 1989), is
labeled. Its colors are consistent with this information, thus placing it higher on the

diagram than a cluster O5 star. Star #82, also labeled, has an unusual placement away

from the main-sequence, probably because it was a blend with the central O star, making

its photometry uncertain. All other stars have colors consistent with cluster membership

and cluster reddening.

Among the 49 stars in the B-V vs. R-I plot are 19 for which Sanders (1972) assigns a

membership probability of >50% from an analysis of the proper motion study by VSA.

There are also eight stars for which Sanders gives a probability of <50%. One of these is

the central O5 star (#80) which, is clearly a cluster member based on symmetry, age,

spectra, and photometry. This star and the other seven have been included in the results

presented here because their colors are consistent with cluster membership.

The V-R vs. R-I diagram (Figure 16) contains data from 127 stars for which V, R, and

I magnitudes were determined, including the 49 stars from the previous plot. Eleven

stars that were blends or had other photometric problems, are marked by ‘x’s and were

excluded from further study. The moderately reddened A1 star (#101) is labeled. Stars

with V < 17.6 (relatively “good” photometry) are shown by filled circles and stars with

V > 17.6 are shown by open circles. The open circles are expected to have errors in V of

0.09 magnitudes or more. As in the previous diagram, the solid curve is the unreddened

main-sequence and the dotted curve is the same line reddened by $E_{(B-V)} = 0.8$. The stars

with good photometry appear to define a heavily reddened main-sequence consistent with

cluster membership.

Unfortunately, in both diagrams, the direction of reddening is nearly parallel to the

slope of the main-sequence curve, meaning foreground stars with moderate reddening of

the correct amount or background stars behind the cluster with very large amounts of
reddening may be confused with the reddened cluster members. However, it is not expected that the density of stars over the 2.4 kpc to the cluster is sufficient to contribute more than a few foreground stars in the small volume defined by the 7' x 7' field of view. In addition, the contribution of background stars lying in the much larger volume behind the cluster is expected to be equally small.

We can estimate the number of background stars likely to be found in the field of view from average galactic star counts. The region of largest extinction in IC 1805 has an $E_{(B-V)} \text{max} = 1.1$ and it is assumed that this defines the total obscuration both within the cluster IC 1805 and in the foreground. By using the tabulation of Allen (1976), a star surface density of 316 per square degree on the sky is appropriate for this galactic latitude ($b = 1\degree$) and to a limiting reddened $V$ magnitude of 17.6 (corresponding to an unreddened $V$ magnitude of $14.1 = 17.6 - R_v[E_{(B-V)\text{max}}]$). This implies that there should only be 4 background field stars apparent within the 7' x 7' field of view to a limiting observed $V$ magnitude of 17.6. Contamination of the cluster sample by such a small number of field stars would have very little effect on the derived mass function. Therefore, all stars except the known foreground A1 star and those represented by 'x's on the color-color plot are assumed to be cluster members.

3.6 THE H-R DIAGRAM

To determine the mass of a star in the cluster, its photometric information needs to be mapped onto the Hertzsprung-Russel (H-R) diagram, a plot of bolometric magnitude ($M_{\text{bol}}$) against the log of the effective temperature ($T_{\text{eff}}$). The placement of the star on
Figure 15. $B-V$ vs. $R-I$ diagram for IC1805. The solid line is the unreddened main sequence taken from Johnson (1966). The dotted line is the main sequence reddened by the average extinction for the cluster, $E(B-V) = 0.8$. The dashed line is the reddening vector showing the magnitude and direction of the average cluster reddening.
Figure 16. $V-R$ vs. $R-I$ diagram for IC 1805. As in fig. 15, the solid line is the unreddened main sequence, the dotted line is the reddened main sequence, and the dashed line is the reddening vector. Stars with $V < 17.6$ are shown as filled circles and stars with $V > 17.6$ are shown as open circles. Open circles have estimated errors in $V$ of 0.07 magnitudes or more. Stars having serious problems with their photometry are shown as 'x's, and were excluded from further study.
this diagram is compared to theoretical evolutionary tracks corresponding to stars of different masses. Color indices were used to estimate both the effective temperatures and the bolometric corrections to the visual magnitudes for each star in the sample. This technique has been adopted previously for the study of OB clusters in the Magellanic Clouds and our own galaxy (Sagar et al. 1986, Massey & Thompson 1991, Massey 1993).

First, observed V-I magnitudes, or B-I for the stars where B magnitudes were available, were dereddened for all assumed cluster members by using a uniform extinction for this region of IC 1805 characterized by an $E_{(B-V)}$ of 0.8. These intrinsic $(B-I)_0$ and $(V-I)_0$ were assigned a main-sequence spectral type using the transformation of Johnson (1966). The spectral type was then transformed to an effective temperature and a bolometric correction (BC) using the listing given in Table 5. The table was assembled from the literature using the temperatures and bolometric corrections of Schmidt-Kaler (1982) for spectral types B through G, and those of Chlebowski & Garmany (1991) for O stars.

Bolometric magnitudes, related to the integrated stellar output of a star over all wavelengths, were derived by adding the bolometric corrections to the intrinsic visual magnitudes after correcting for the cluster distance using the distance modulus (DM) to IC 1805 of 11.9 (G&V, 1989). The intrinsic visual magnitudes are equal to the observed visual magnitudes corrected for interstellar absorption. Thus, the formula is:

$$M_{bol} = (V - R_V * E_{(B-V)}) - DM + BC$$

(7)

Figure 17 shows the H-R diagram for the data. The average uncertainty in the location of any star on this plot, attributable to carrying forward the estimated photometric uncertainties, is shown in the lower right corner. Included in the figure are post-main-
Table 5. Table showing calibration of color indices versus spectral type, effective temperature and bolometric correction. Compiled from Johnson (1966), Schmidt-Kaler (1982) and Chlebowski & Garmany (1991).

<table>
<thead>
<tr>
<th>(B-I)₀</th>
<th>(V-I)₀</th>
<th>Spectral Type</th>
<th>T eff</th>
<th>BC</th>
<th>(B-I)₀</th>
<th>(V-I)₀</th>
<th>Spectral Type</th>
<th>T eff</th>
<th>BC</th>
</tr>
</thead>
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<td>--</td>
<td>--</td>
<td>O4</td>
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<td>--</td>
<td>--</td>
<td>A1</td>
<td>9230</td>
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</tr>
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<td>O5</td>
<td>44300</td>
<td>-4.06</td>
<td>0.15</td>
<td>0.09</td>
<td>A2</td>
<td>8970</td>
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</tr>
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<td>-0.79</td>
<td>-0.47</td>
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<td></td>
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<td>8720</td>
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</tr>
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<td>42200</td>
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<td>0.36</td>
<td>0.22</td>
<td>A5</td>
<td>8200</td>
<td>-0.15</td>
</tr>
<tr>
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</tr>
<tr>
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<td>-0.47</td>
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<td>40100</td>
<td>-3.77</td>
<td></td>
<td></td>
<td>A8</td>
<td>7580</td>
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</tr>
<tr>
<td>--</td>
<td>--</td>
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<td>39100</td>
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<td>0.78</td>
<td>0.47</td>
<td>F0</td>
<td>7200</td>
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</tr>
<tr>
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<td>6890</td>
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</tr>
<tr>
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<td>0.64</td>
<td>F5</td>
<td>6440</td>
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<td>6200</td>
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</tr>
<tr>
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<td>0.81</td>
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<td>6030</td>
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<tr>
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<td>30000</td>
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<td>5860</td>
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<td>1.55</td>
<td>0.89</td>
<td>G5</td>
<td>5770</td>
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<tr>
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<td>-0.32</td>
<td>B2</td>
<td>22000</td>
<td>-2.35</td>
<td>1.70</td>
<td>0.96</td>
<td>G8</td>
<td>5570</td>
<td>*</td>
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<tr>
<td>-0.47</td>
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<td>18700</td>
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<td>1.88</td>
<td>1.06</td>
<td>K0</td>
<td>5250</td>
<td>-0.31</td>
</tr>
<tr>
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<td>-0.22</td>
<td>B5</td>
<td>15400</td>
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<td>--</td>
<td>--</td>
<td>K1</td>
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<td>-0.37</td>
</tr>
<tr>
<td>-0.33</td>
<td>-0.19</td>
<td>B6</td>
<td>14000</td>
<td>-1.21</td>
<td>2.14</td>
<td>1.22</td>
<td>K2</td>
<td>4900</td>
<td>-0.42</td>
</tr>
<tr>
<td>-0.29</td>
<td>-0.17</td>
<td>B7</td>
<td>13000</td>
<td>-1.02</td>
<td>--</td>
<td>--</td>
<td>K3</td>
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<td>-0.50</td>
</tr>
<tr>
<td>-0.21</td>
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<td>11900</td>
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<td>--</td>
<td>--</td>
<td>K4</td>
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</tr>
<tr>
<td>-0.12</td>
<td>-0.06</td>
<td>B9</td>
<td>10500</td>
<td>-0.51</td>
<td>2.77</td>
<td>1.62</td>
<td>K5</td>
<td>4350</td>
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</tr>
<tr>
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<td>0.00</td>
<td>A0</td>
<td>9520</td>
<td>-0.30</td>
<td></td>
<td></td>
<td>K7</td>
<td>4060</td>
<td>-1.01</td>
</tr>
</tbody>
</table>

* Anomalous B.C. value of -0.40 given by Schmidt-Kaler.

sequence evolutionary tracks taken from Maeder and Meynet (1988) for O7 to B5 stars, and pre-main-sequence tracks from Iben (1965), and Cohen and Kuhi (1979) for B5 to G0 stars. Using post-main-sequence tracks for all masses would change the mass spectrum only slightly, but would lead to the conclusion that many of the low-mass stars
Figure 17. The H–R diagram for IC 1805. Filled circles represent the data for \( V < 17.8 \) and open circles for \( V > 17.8 \). Dashed lines are post–main sequence evolutionary tracks taken from Maeder & Meynet (1988). Solid lines are pre–main sequence tracks taken from Cohen & Kuhi (1979) with the exception of the 6 solar mass track which is from Iben (1965). Dotted lines labeled A–H are the isochrones corresponding to these tracks.
were $10^8$ to $10^9$ years old. Since there are no evolved stars of higher mass, it is safe to assume that these stars are young and still evolving toward the main-sequence. Pre-main-sequence isochrones are also shown indicating stellar ages ranging from $10^5$ - $10^7$ yr., similar to the age quoted by Sagar et al. (1986) in their study of IC 1805. This finding implies that there was no disruption of star formation $3\times10^6$ years ago as suggested by G&V (1989). It must be noted, however, that unresolved binaries with inflated magnitudes, if present in the data, would have overestimated masses and underestimated ages, thus distorting the mass spectrum and pushing down the lower age limit.

3.7 THE MASS FUNCTION

According to Scalo (1986) the mass spectrum $f(m)$ is defined as the number of stars born per unit mass interval $dm$ while the mass function $F(\log m)$ is the number of stars born per unit logarithmic (base ten) mass interval $d\log m$. Normalization of $F(\log m)$ by the area of sky under observation (in square kiloparsecs) gives the "initial mass function" $\xi(\log m)$, as long as all stars under study are assumed to be of similar age. The parameter of frequent interest is the logarithmic slope of $\xi(\log m)$,

$$\Gamma(m) = \frac{d \log \xi(\log m)}{d \log m}$$

(8)

which, for a power-law mass spectrum, is independent of mass.

From the H-R diagram (Figure 17), the number of stars within each mass bin can be determined. Dividing by the log mass range of each bin and the effective area observed ($2.39 \times 10^{-5}$ kpc$^2$) the initial mass function $\xi(\log m)$ can be determined. Table 6 presents a summary of the results for each mass bin and Figure 18 shows $\log \xi$ plotted against the
Figure 18. The mass function for IC 1805. Filled circles represent the data from this study when only stars with $V < 17.6$ are considered. Open circles represent the two lowest mass bins when all stars are considered. The solid line is a weighted least-squares fit to the filled circles (excluding the lowest mass bin) yielding a slope of $\Gamma = -1.28 \pm 0.31$. Triangles represent data from Sagar et al. (1986) with a weighted fit (dotted line) excluding the two lowest mass bins giving $\Gamma = -1.25 \pm 0.17$. Error bars show root(\textit{N}) deviations.
log of the mean mass. Filled circles represent data from this study having “good” photometry ($V < 17.6$), with error bars showing $\sqrt{N}$ deviation where $N$ is the number of stars in the mass bin. The open circles indicate the values for $\log \xi$ when all stars are considered. The solid line in Figure 18 is a least-squares fit to the open circles, weighted by the number of stars in each bin and excluding the lowest mass bin (due to probable incompleteness). The line has a slope of $-1.28 \pm 0.31$, which is very close to the value of $-1.35$ for the field star IMF as determined by Salpeter (1955).

Table 6. IC 1805 Mass Function

<table>
<thead>
<tr>
<th>Mass Range</th>
<th>log($M/M_\odot$)</th>
<th>N ($V&lt;17.6$)</th>
<th>log($\xi$) ($V&lt;17.6$)</th>
<th>N all stars</th>
<th>log($\xi$) all stars</th>
</tr>
</thead>
<tbody>
<tr>
<td>30 - 15 M$_\odot$</td>
<td>1.33</td>
<td>2</td>
<td>5.44</td>
<td>2</td>
<td>5.44</td>
</tr>
<tr>
<td>15 - 9</td>
<td>1.07</td>
<td>6</td>
<td>6.05</td>
<td>6</td>
<td>6.05</td>
</tr>
<tr>
<td>9 - 5</td>
<td>0.827</td>
<td>6</td>
<td>5.99</td>
<td>6</td>
<td>5.99</td>
</tr>
<tr>
<td>5 - 3</td>
<td>0.588</td>
<td>23</td>
<td>6.64</td>
<td>23</td>
<td>6.64</td>
</tr>
<tr>
<td>3 - 2.25</td>
<td>0.415</td>
<td>14</td>
<td>6.67</td>
<td>14</td>
<td>6.67</td>
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<tr>
<td>2.25 - 1.5</td>
<td>0.264</td>
<td>35</td>
<td>6.92</td>
<td>40</td>
<td>6.98</td>
</tr>
<tr>
<td>1.5 - 1.25</td>
<td>0.137</td>
<td>3</td>
<td>6.20</td>
<td>23</td>
<td>7.08</td>
</tr>
</tbody>
</table>

Sagar et al. (1986) performed a photoelectric UBV survey of stars whose membership probability exceeds 50% based on the proper motion study by VSA (1965) and the revision by Sanders (1972). The survey covered a region of IC 1805 of approximately 3000 arcmin$^2$ centered near ADS 1920 which is far larger than the 49 arcmin$^2$ region of this study. Sagar et al. (1986) transformed these measurements onto the H-R diagram and determined the mass function $F(\log m)$ for a total of 136 stars. They quote a logarithmic slope for the mass function of $-1.43 \pm 0.27$ for tracks that assume no mass loss. Their data are included in Figure 18 for comparison (solid triangles). It was necessary to normalize the data from their table 3 by the width of the mass bin (0.1) and
mass bin (0.1) and the area surveyed (3000 arcmin$^2 = 1.46 \times 10^{-3}$ kpc$^2$). The dashed line is a least-squares fit to the data weighted by the number in each bin and excluding the two lowest mass bins. The slope of the line is $-1.25 \pm 0.17$.

The slopes for the two mass functions are the same within the uncertainties, thus implying that only the absolute rate of star formation is different between the inner and outer regions of IC 1805 for stars in the overlapping mass range. The earlier study shows a turnover at log($M/M_\odot$) = 0.6, which is attributed to incomplete measurements of fainter stars. The present study proves this assertion correct as it shows no turnover until log($M/ M_\odot$) = 0.2. It is assumed that the present study also suffers from incompleteness at the bottom of the mass range.

Another galactic OB association Cyg OB2 was studied by Massey & Thompson (1991) to determine its mass function. Cyg OB2 is a highly obscured cluster [$E_{(B-V)} = 1.3$] containing one of the few classified O3If stars and the ultra-luminous B supergiant VI Cyg OB#12. Massey & Thompson’s study of a $3.13 \times 10^{-4}$ kpc$^2$ region centered near SAO 49775 results in a mass function with a slope of $\Gamma = -0.95 \pm 0.23$, shown in Figure 19, when their foreground subtracted data are considered (triangles and dashed line). The four top mass bins of this study are included for comparison (open circles). The absolute values of the two mass functions are surprisingly similar, indicating similar rates of star formation. In the region of overlap, the slopes of the two mass functions also appear to be similar, but it must be noted that the errors in the high mass data points shown for IC 1805 (Figure 18) are large due to few stars.
Figure 19. The mass function for Cyg OB2 compared to the top mass bins from this study. The solid line is the same as in fig. 18. The dotted line is a weighted least-squares fit to the triangles (excluding the lowest mass bin) giving a slope of \( \Gamma = -0.95 \pm 0.23 \). Error bars are root(N) deviations.
4 CONCLUSIONS

Large-format CCDs make excellent tools for conducting photometric studies of open clusters, but proper care must be taken when reducing the data to ensure accurate calibration of the output and to lessen the amount of noise introduced during this reduction process. Observing must be carried out with an understanding of noise contributions to balance time and storage resources optimally between target images and calibration frames. Because the literature is far less complete for R and I passbands than for U, B, and V, observation of nearby standard stars is recommended for calibrating R and I magnitudes in open clusters.

The slope of the mass function for the central region of IC1805 was found to be -1.28 ± 0.31 for masses > 1.5 M☉, similar to the slope of the field star IMF and similar to the slope found for a previous study of the most massive members of the entire cluster. The absolute value of the mass function appears to be larger for the central region than for the cluster as a whole, indicating a higher rate of star formation there. Also, evidence suggests that star formation in the cluster has occurred over ten million years, making the observed mass function an average of the initial mass function over this time.
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