STELLAR PHOTOMETRY WITH AN AUTOMATED MEASURING MACHINE

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ABSTRACT

The combination of wide field high quality telescopes, fine grain emulsions and fast automated measuring machines offers an unrivalled opportunity for progress in statistical astronomy. This review illustrates, with examples, the many steps which are necessary to realise this potential.

1. INTRODUCTION

The availability of large numbers of deep high quality photographic plates covering wide fields, and containing perhaps 100,000 usable images, presents new challenges in data reduction and data handling and new opportunities for statistical studies in astronomy. The existence of new fast automated measuring machines (COSMOS and APM in the UK) provides an unrivalled opportunity to exploit these data. Such exploitation however requires a careful understanding of the pitfalls as well as the advantages of automated photometry. This review outlines the practical factors involved in converting a Schmidt telescope plate into quantitative science, using as an example a survey of the stellar distribution in the Galactic spheroid. While most of the specific examples are relevant to plates from the UK Schmidt telescope measured with the COSMOS machine in Edinburgh, the principles are more generally valid. This paper discusses in turn the following topics: the operation of the measuring machine, and the significance of its parameters (section 2); the matching of several plates onto a single master reference set and image classification (section 3); photometric calibration (section 4); and some examples of the potential of fast measuring machines and Schmidt plates in statistical astronomy (section 5). Many other examples are presented in other contributions to this volume.

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2) THE MEASURING MACHINE or WHAT HAS THAT MACHINE DONE TO MY PLATE?

The automated measuring machines may be used in either of two ways. In the first, they operate exactly as any other microdensitometer, recording the coordinate and density of every pixel in an area of plate. In the second, a real time decision is made by the operating system as to whether a pixel is above or below some prespecified threshold relative to the local sky background. Those pixels above this threshold are recorded, while those below are used to improve the current estimate of the local sky value. This procedure vastly reduces the number of pixels which need to be recorded, from the $\sim 10^9$ which are required to map all information on a single large Schmidt plate, to more manageable numbers. Apart from saving disk storage space, this greatly increases the speed of the machines, allowing an entire UK Schmidt plate to be mapped with 30μ pixels in less than half a day. After this measurement, the recorded pixel densities are converted to intensities, using an appropriate calibration curve, and processed by an algorithm which looks for connected pixels, joins them into an image, and deduces some coordinate, shape, structural, and total intensity parameters. It is these which are presented to the unsuspecting astronomer, and the derivation of these which must be understood to allow useful analysis of the results.

There are several parameters which must be specified for this stage of measurement which have a significant effect on the resulting data. These include the size of the measuring spot (the pixel size), the step size between consecutive measurements, the density to intensity conversion, the scale length for updating the local sky background value, the threshold above this level at which to accept pixels as 'signal', and the minimum number of joined pixels which constitute a 'real' image.

a) Pixel size: This is normally 16μ or 32μ with COSMOS and 7μ with the APM, although, as the intensity profile of the measuring spot shows extended wing structure, the physical significance of these numbers is unclear. In practice, the shape of the spot is a significant feature only in areas of steep density gradients.

b) Pixel spacing: This may usually be set at one-half the pixel size, following the Nyquist criterion. In practice, the optimum pixel size and spacing is dependent on the grain structure of the emulsion, the telescope plate scale, and the magnitude of the images of interest. A detailed analysis of these parameters on real plates has been carried out by Okamura et al. (1983), where details may be found.

c) Determination of sky: The primary disadvantage of fast measuring machines is their speed. This forces the use of relatively simple algorithms for sky estimation, and consequent high thresholds. The most reliable way to determine the sky background on a plate is

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obviously to scan the whole plate, fit some suitable polynomial to the data to estimate the background, and then rescan the plate. This however presupposes that the machine retains its coordinate and intensity system, and doubles the machine time per plate. A second method is to smoothly extrapolate the expected sky value from the values measured to date. A further complication is the usual choice of mean or mode as the best estimator of the true value. The practical consequence of these points is that typical thresholds above sky are 5%-10% - very high by the standards of those doing PDS photometry of galaxies - and that significant systematic errors in photometry, correlated with the local 'sky' value, can arise (cf section 4). The most important effect of these thresholds is to systematically distort the magnitude scale at faint magnitudes. This is discussed further in section 4.

A substantial improvement in this situation is possible, at the expense of considerable computer storage and time, by digitising and storing the entire plate. This then allows more sophisticated algorithms to be utilised to correct for variations in sky background, telescope vignetting and geometrical distortion, and so on. This technique is regularly applied to prime focus 4m plates, and allows thresholds of 1%-2% of sky to be reached, in spite of the extreme variations across these plates. See the COSMOS Users Guide for further details.

d) The density to intensity conversion: The choice of a calibration curve which is appropriate to a specific pixel at a specific position is an extremely complex problem. Factors which must be considered include the choice of wedge (continuously variable) or (discrete) spot calibration, the effect of exposing the calibration on sky or clear plate, the effects of time delays between image and calibration exposures and plate development, real variations in the true density-intensity relation with position and/or wavelength, and the systematic consequences of random errors in the adopted relation. These and other factors have been recently reviewed elsewhere (Gilmore 1983a) and so will not be repeated here. We note however that in practice it is not usually possible to do more than adopt a spline fit to the calibration provided on the plate. Projects which require a higher precision calibration must accept that very considerable effort is required to provide it.

e) Minimum image size: The minimum number of pixels in a 'real' image is a complex function of the seeing during the exposure, the telescope and emulsion scattering properties, the machine spot intensity profile, and the threshold set during measurement. In real cases, the minimum size is best set too low, so that most of the smallest (faintest) images are spurious. These may then be rejected by visual comparison with the original, or during the comparison with another plate (see below). While this is relatively straightforward, it is a very important consideration if faint object number counts are to be used as a test of completeness. It is not impossible to find a combination of too low a threshold and too small a minimum image size which mimics plausible number-magnitude counts much fainter than the true plate "limiting" magnitude. This problem is best eliminated, as are most others, by using several plates in each of several colours, and retaining images common to all (see below) and by remembering to look at the images on a plate which correspond to the machine output.

3) INITIAL PROCESSING *or* WHAT DO I DO NOW?

After the machine processing outlined above, an astronomer will be presented with a set of magnetic tapes containing information on the position, intensity and shape of each of perhaps 10^5 groups of pixels identified by the algorithm as an 'image'. Many will be spurious. The next stage of processing is best illustrated assuming that serious photometry is intended. In this case, several plates in perhaps several colours on the same field centre have been measured. It is then necessary to transform them all to the same coordinate grid, match 'real' images, and classify them.

a) Merging: This involves the transformation of the coordinate frame of each plate onto that of some master plate. In practice, this is achieved using a grid of stars common to all plates, which are identified (visually or by cross-correlation) and used to define the usual multi plate-constant transformation. For plates from the same telescope and with the same field centre such a transformation is usually accurate globally to a few microns, or ~ 0.1 pixels, allowing transformations to an astronomical reference frame with an accuracy of a few tenths of an arcsecond. The stars chosen for this initial transformation should be neither very faint (to minimise random errors in position) nor very bright (to minimise systematic errors due to the inclusion of asymetric image halo structure).

b) Matching: When all image centroids are defined on the same reference frame, which may be an RA-DEC grid, it is straightforward to identify images which appear on more than one plate simply by position coincidence. This simple process however largely determines what science can be done with the resulting data, and so must be carefully considered. Some examples will illustrate this point. A common procedure is to take the 'deepest' plate available, and match all others onto it. This plate in practice is likely to be a sky limited IIIa-J plate, reaching to B ~ 22.5 or 23.0. The matching of an I band plate (limiting magnitude ~ 19.5) to this J plate will obviously retain every image on the I plate with B-I colour less than ~ 3^m. Thus, while the vast majority of images will be retained, a systematic rejection of late M dwarfs will result. This rejection is so efficient that the inverse procedure has been used with considerable success to identify extremely red stars (Gilmore and Hewett, 1984). Other classes of object which may be strongly biased against include variables and proper motion stars.

A more important effect for non-moving objects of neutral colour - just those usually required - is the effect of unresolved images. Deep Schmidt plates are already resolution limited over most of the sky (thereby precluding deeper surveys), so that a significant number of images are either unresolved or marginally resolved from other images. When adjacent images have different colours, or different plates are taken in different seeing, many images can be resolved on some and unresolved on other plates of a set. The reliable handling of these images is a very major task when processing large data sets, but must be carefully considered. At galactic latitude 50° below the Galactic bulge, 10% of images are affected by merging independent of apparent magnitude. These images must be reliably identified and handled in constructing the final data set. They are best found by comparison with deeper, larger scale plates. [The detailed analysis of crowded field photometry will not be discussed here. An excellent description is given by Newell (1983).] In the determination of the number of unresolved image pairs, it is striking how rapidly the surface density of stars may change across a Schmidt plate due to the variation in Galactic coordinates over the 6° field. Corrections for image crowding should be carried out in galactic rather than rectilinear plate coordinates. These corrections are themselves a useful scientific result, as they nicely map the gradient in the stellar surface density of the Galaxy. The determination of this is often the aim of the original project.

c) Image classification: When a final merged and matched set of images has been produced, it is usually necessary to decide what is a star and what is a galaxy. Images may be classified into one of four obvious categories:- definite star; definitely resolved; definite multiple image; and don't know. This classification can be carried out using surface brightness, image shape or extent, colour and other parameters. Examples are given by Kron (1980) and Reid and Gilmore (1982), and references in those papers. The details need not be discussed here. The important consequence for the intended scientific use is that no single classification scheme is optimised for all projects. A project which must include all stellar objects can do so only by accepting some galaxy contamination when classifying uncertain images, and vice versa. Provided that the consequences of this choice are known, and adequately calibrated this is not a problem.

An example of such a post-classification check is shown in Figure 1, which is the two-point correlation function for galaxies and faint K stars towards the south Galactic pole. The absence of any structure in the 'stellar' distribution on scales at which it occurs in the 'galaxy' distribution, at the same apparent magnitude and colour limits on the same plate, is a very powerful diagnostic of possible galaxy contamination of the stellar sample. [The lack of any significant structure in the stellar distribution is also a useful proof that no patchy obscuration is present in this field. cf Gilmore et al. 1984 for details.]

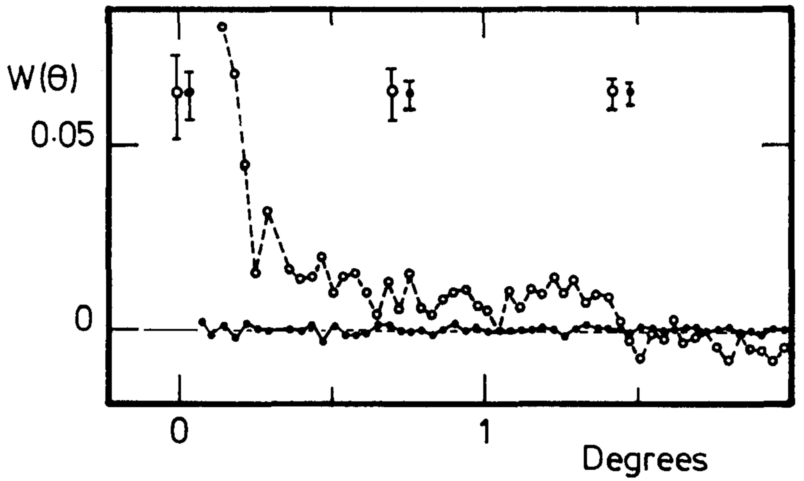

Figure 1 - the 2-point correlation function for galaxies (open points) and stars (solid points) at the south galactic pole. The absence of a galaxy feature in the stellar distribution precludes galaxian contamination of the stellar sample, and patchy interstellar reddening.

After these steps, the astronomer has a magnetic tape containing a coordinate, a shape, a label (star, resolved, etc.), and some intensity parameter for each of (possibly) several measures of every reliable image in some master list. It is now time to attempt photometry.

4) PHOTOMETRIC CALIBRATION or HOW BRIGHT IS THAT STAR?

Photographic photometric surveys divide naturally into those which require a reliable differential measure of luminosity or colour, and those which require 'true' magnitudes, free from zero point or colour scale errors.

a) Uncalibrated Photometry: An example of a project using purely differential photometry is a study of variable objects, where the aim is to find an image whose ranking, relative to a suitable set of nearby images, has changed significantly. The general problem of the detection of variability from uncalibrated data has been discussed by Gilmore (1979) and applied by Hawkins (this volume) to the particular case of faint quasars. A further discussion of this is given elsewhere (Gilmore, 1983a), and will not be repeated here.

b) Magnitude Scale Calibration: The basic philosophy underlying the calibration of magnitude scales using photometric standards which span the position, colours and magnitudes of interest is the application of consistent systematic errors. As the measurements of the standard are in error by exactly the same amount as the other

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images of interest, their use to determine the magnitude scale and zero point will automatically eliminate all these errors. A practical example of the application of this philosophy is described in Reid & Gilmore (1982), and the general problem is discussed by Gilmore (1983a). The following important points are discussed below: the meaning of a magnitude scale; the limits on systematic errors, the establishment of faint standards; and the accuracies attained in practice.

i) Types of magnitude: Apparent magnitudes for the relatively bright stars used as secondary standards in most photometry are usually obtained by photoelectric aperture photometry. Apparent magnitudes for faint standards are typically measured from sub-beam (Pickering-Racine) prism plates, or from CCD frames. These may produce magnitudes which are defined by either an aperture or a threshold or which may be total. There is no particular reason to expect these magnitude scales to agree either with each other, or with other observers. Considerable care must be taken when establishing faint standards that their magnitudes are defined in the same way as those of brighter stars, and in a way which is appropriate for their intended use. An illustration of the effect of thresholding on total magnitudes, of errors in the adopted local sky background, and of a real change in the local sky background (due to unresolved stars or gradients in the zodiacal light, for example) is shown in Table 1 (P.C. Hewett, pri.comm'n).

Table 1

Total Magnitude	10% Thresholded Magnitude	Sky error of +1%	Sky change of +3%
19.50	19.69	19.72	19.73
20.50	20.87	20.91	20.92
21.50	22.33	22.43	22.41
22.00	23.42	23.66	23.76
22.40	25.35	28.06	-

Calculated by P. Hewett using a sky brightness of 22.5 mag/arcsec^2, and a Moffat stellar profile with R = 2.5 and B = 3.0

ii) Establishing Faint Standards: Provided the points above are taken into consideration, reliable faint magnitudes can be established using sub-beam prism and CCD data. Recently Blanco (1982) has questioned the reliability of sub-beam prisms. Examples of the successful use of these devices, provided adequate care is taken, are presented by Reid & Gilmore (1982), Christian & Racine (1983) and Gilmore (1983a, esp. Figure 7). As always however, the only reliable test of a faint sequence is an independent attempt to derive a magnitude scale using several methods. For the study of the stellar distribution in the Galactic spheroid described by Gilmore (1983b), this is achieved by using photoelectric aperture photometry, sub-beam prism photographic photometry, and CCD photometry from two separate cameras in each field of interest. While this calibration then involves an enormous workload, and considerable overlap of observations, it does guarantee that the calibrating standards are themselves reliable - a sine qua non for useful photometry.

iii) Systematic errors: When a reliable consistent set of standards have been derived, they will usually be found inadequate to define any structure in the relation between instrumental and standard magnitudes. Some structure in the form of a change of slope of the relation is to be expected near the apparent magnitude at which the plate and measuring machine becomes saturated ($m \sim 19$) and at the magnitude when the image halo rises above the threshold, and becomes included in the total image ($m \sim 12$). An example of this is seen in Figure 2, which also illustrates the excellent agreement between photoelectric aperture and CCD magnitudes and those derived from sub-beam prism photometry. For stars with $V \leqq 11^m$, a systematic error of up to $0^m\!.2$ may be present, while the magnitude scale is undefined for $V > 19.5$. As much published photometry is based on substantially fewer standard stars per magnitude interval than the example shown here, significant non-linearities in that data are possible.

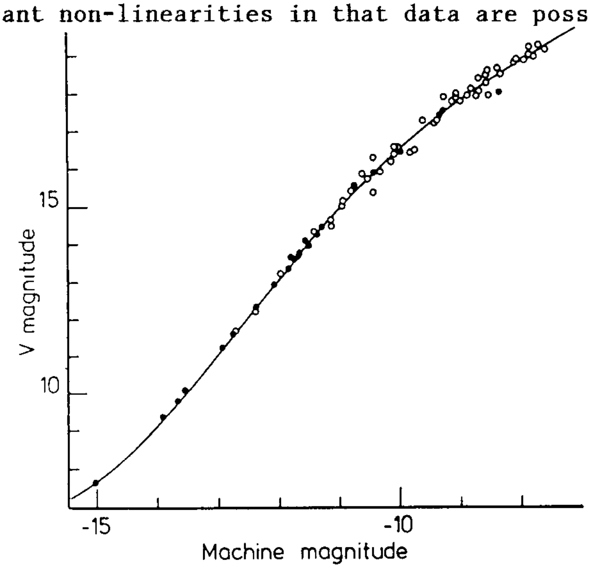

Figure 2 - A Schmidt plate calibration curve using standards from photoelectric and CCD observations (open points and all with $V < 12$) and sub-beam prism plates (solid points). The curvature is evident.

The second major class of systematic errors includes all those effects where a correlation exists between the residual magnitude from the calibration curve and a second parameter. The best known are those correlations with position ("field effects" - see Reid & Gilmore

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(1982) for a discussion), with colour ("colour equations", cf Faber et al. (1976), and Blair and Gilmore (1982)) and with local sky background (cf Table 1). The existence of these effects must be carefully tested if reliable photometry is required.

iv) Random errors: If all the tests and calibrations discussed above have been carried out, a final object list with photometry is available. How good is it? A reliable estimate of the random errors can be derived by independently calibrating each plate in a set, and calculating the rms dispersion in the derived magnitudes for every star in the sample. Table 2 shows such results for a sample of ~ 20,000 stars with photometry from 5 V, 3 B and 3 I plates. This table shows that, with care and a lot of effort, reliable magnitudes and colours for large samples of stars can be derived with random errors of a few percent.

Magnitude	RMS(V)	RMS(B)	RMS(I)
10.5	0.09	0.07	0.07
11.5	0.06	0.09	0.08
12.5	0.05	0.06	0.10
13.5	0.05	0.06	0.10
14.5	0.05	0.06	0.10
15.5	0.06	0.07	0.09
16.5	0.06	0.07	0.08
17.5	0.08	0.07	0.11
18.5	0.15	0.09	0.21
19.5	0.26	0.19	-
20.5	0.42	0.34	-

5) SOME RESULTS or WHY DID I DO THAT?

What does one do with 20,000 accurate magnitudes and colours? This volume contains many examples of the type of astronomy which can be done only with large samples of carefully measured objects. Two others are shown in Figures 4 and 5 below.

Figure 3 shows the U-B/B-V diagram for a field in the southern hemisphere which is part of a survey for objects which do not have main sequence stellar colours. Some 250 square degrees have been surveyed as a pilot project, to prove the potential of this method for finding white dwarfs, HB stars, cataclysmic variables, and bright quasars. The high precision of the photometry and the very large number of objects measured allow a careful discussion of completeness and selection effects, and the detection of objects with less extreme colours than most other similar surveys to date. In particular, cooler white dwarfs and higher redshift quasars can be found. Further details are in Gilmore (1983c).

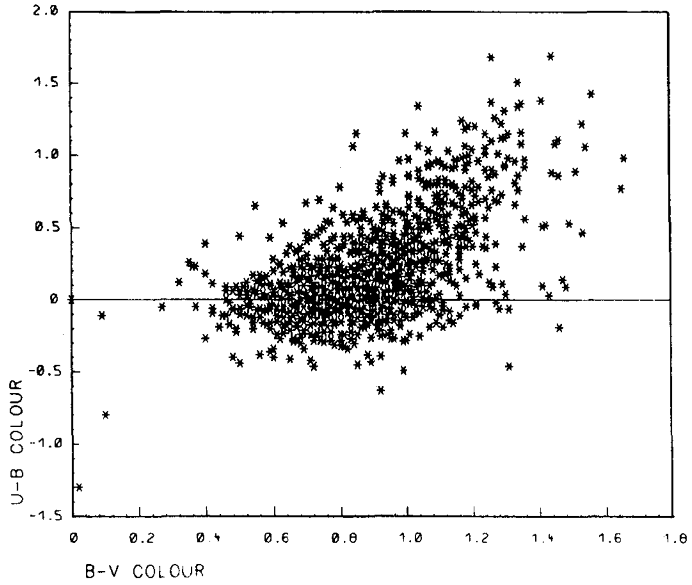

Figure 3 - A 2-colour diagram from a survey for stellar objects with non main-sequence colours.

Figure 4 shows the stellar number-magnitude-colour distribution towards the south Galactic pole. Similar data have been obtained in nine fields, and are being analysed to determine the structure and evolution of the Galactic spheroid, and the solar neighbourhood luminosity function. Further details may be found in Gilmore (1983b, 1983d, 1984) and Gilmore & Reid (1983).

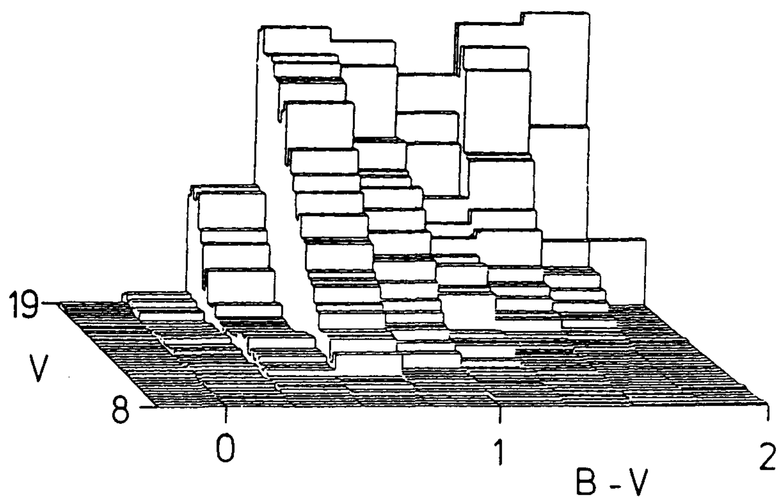

Figure 4 - The stellar number-magnitude-colour distribution of 22,000 stars near b = -90°. The prominent bimodality is evident for V > 17.

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DISCUSSION

V.M. BLANCO: I must remark that the use of the Pickering prisms or so-called Pickering-Racine wedges can result in large errors in the extrapolation of photometric sequences. This led us to stop their use at Cerro Tololo. The main problem is that the beam produced by the small prism has a different f/ratio than the beam produced by the telescopes' full aperture. This results in appreciably different density structures for the primary and secondary images of a given star. So any of you who desire to use such a prism or wedge I recommend a careful reading of Racine's own paper wherein he reviewed this old idea. Also, you may want to read my paper in the PASP where are discussed the sizeable errors that can result from the use of these prisms.