The colour–magnitude relation for galaxies in the Coma cluster

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ABSTRACT

We present a new photometric catalogue of the Coma galaxy cluster in the Johnson U and V bands. We cover an area of 3360 arcmin^2 of sky, to a depth of V^\hat{ } = 20 mag in a 13-arcsec diameter aperture, and produce magnitudes for 1400 extended objects in metric apertures from 8.8- to 26-arcsec diameters. The mean internal rms scatter in the photometry is 0.014 mag in V, and 0.026 mag in U, for V_{13} < 17 mag.

We place new limits on the levels of scatter in the colour–magnitude relation (CMR) in the Coma cluster, and investigate how the slope and scatter of the CMR depend on galaxy morphology, luminosity and position within the cluster. As expected, the lowest levels of scatter are found in the elliptical galaxies, while the late-type galaxies have the highest numbers of galaxies bluewards of the CMR. We investigate whether the slope of the CMR is an artefact of colour gradients within galaxies, and show that it persists when the colours are measured within a diameter that scales with galaxy size. Looking at the environmental dependence of the CMR, we find a trend of systematically bluer galaxy colours with increasing projected radius from the centre of the cluster. Surprisingly, this is accompanied by a decreased scatter of the CMR. We investigate whether this gradient could be caused by dust in the cluster potential, however the reddening required would produce too large a scatter in the colours of the central galaxies. The gradient appears to be better reproduced by a gradient in the mean galactic ages with projected radius.

Key words: catalogues – galaxies: clusters: individual: Coma – galaxies: fundamental parameters.

1 INTRODUCTION

The progressive reddening of the integrated colours of elliptical galaxies with increasing luminosity is known as the colour–magnitude relation (CMR) (Faber 1973; Visvanathan & Sandage 1977; Frogel et al. 1978; Persson, Frogel & Aaronson 1979; Bower, Lucey & Ellis 1992a,b, hereafter BLE92a, BLE92b). Despite being a far simpler relation than the Fundamental Plane (Dressler et al. 1987; Djorgovski & Davis 1987; Bender et al. 1992; Saglia, Bender & Dressler 1993; Jorgensen, Franx & Kjærgaard 1993; Pahe, Djorgovski & De Carvalho 1995), it none the less has similarly small levels of scatter. Traditionally, the slope seen in the CMR has been attributed to a mass–metallicity sequence (Dressler 1984; Vader 1986), with the massive galaxies being more metal-rich, and thus redder, than the less massive ones. This tendency can naturally be explained by a supernova-driven wind model (Larson 1974; Arimoto & Yoshii 1987), in which more massive galaxies can retain their supernova ejecta for longer than can smaller galaxies, thus being able to process a larger fraction of their gas before it is expelled from the galaxy. In hierarchical models of galaxy formation (e.g. Kauffmann & Charlot 1998), the CMR can be reproduced because metals are expelled from low-mass galaxies as part of the feedback process. Studies of the CMR in high-redshift clusters find a ridge-line slope comparable to that of the local clusters; furthermore, there is no sign of a change in the range of magnitude over which the CMR may be traced (Ellis et al. 1997; Stanford, Eisenhardt & Dickinson 1998; Kodama & Arimoto 1997; Kodama 1997; Kodama et al. 1998). These studies all point towards the galaxies which make up the CMR in the cores of rich clusters being primarily constituted of uniformly old stellar populations. Given this metallicity-driven interpretation of the CMR, its low levels of scatter in cluster cores implies that the galaxies are made up from uniformly old stellar populations (BLE92b; Bower, Kodama & Terlevich 1998). Even small variations in the ages of the galaxies would lead to unacceptable levels of scatter in young stellar populations, whereas old stellar populations have a much smaller age dependency in their colours.

Despite the uniformity of the CMR between the Coma and Virgo cluster cores (BLE92b), studies of the CMR in Hickson compact groups (Zepf, Whitmore & Levinson 1991) show increased scatter.
Similarly, studies of field ellipticals (Larson, Tinsley & Caldwell 1980) indicate that the scatter in the CMR could depend on environment. However these group and field galaxy samples contain data from many disparate sources, therefore there might be an added source of scatter from matching the various photometric data sets on to a single photometric system. These additional sources of scatter can take the form of uncertainties in $K$ corrections for galaxies at different redshifts, or biases introduced as a result of sampling only the brighter end of the luminosity function at higher redshift in a magnitude-limited sample. Despite this, at least in the sample of Larson et al. (1980), the extra scatter seems too large to be accounted for by increased observational uncertainties alone. Kodama et al. (1999) analysed the CMR in the Hubble Deep Field North. Again they found an increase in the CMR scatter, and a possible indication that the slope of the CMR is flatter at high redshift in the field environment.

Further evidence for an environmental dependence of the CMR comes from studies of spectral line indices. Broad-band colours are notoriously inefficient at separating the effects of age and metallicity on a stellar population, a degeneracy neatly summarized by Worthey (1994) in his ‘2/3’ law $[\Delta [\text{Fe/H}] \sim \Delta \log(t)]$. This has led to studies of the stellar populations of early-type galaxies in the cores of clusters using spectral line indices chosen to break this degeneracy and disentangle the effects of age and metallicity. Such studies (e.g. Mehlert et al. 1998; Kuntschner 1998; Kuntschner & Davies 1998) show that the CMR is driven by metallicity variations with galaxy luminosity, rather than age. A direct comparison of colours and line-index methods by Terlevich et al. (1999) showed that the two approaches have comparable sensitivity.

The key question then is to determine how the uniform CMR seen in cluster cores transforms into the less tight CMR of field galaxies. In order to investigate how the CMR varies across the Coma cluster, between different regions and between different galaxy morphological types, we have undertaken a survey of $(U - V)$ colours covering almost 1 deg$^2$ of the Coma cluster. Although the colours of the galaxies in Coma have been studied before, both in a wide area (e.g. Dressler 1980; Godwin, Metcalfe & Peach 1983, hereafter GMP) and with high precision $U$- and $V$-band CCD data (BLE92a), the present study is unmatched in area and sensitivity to variations in stellar populations. We chose to use the Johnson (Johnson & Morgan 1953) $U$ and $V$ filters because they straddle the 4000-Å break in the spectra of galaxies at low redshift, and are thus very sensitive to the ages of the stellar populations. They are especially sensitive to recent bursts of star formation (e.g. Worthey 1994; Charlot & Silk 1994).

The present study extends the photometry of BLE92a to a complete galaxy sample covering approximately four times the area (still centred on the core), and reaches to fainter limiting magnitudes. The extra coverage and depth will enable us to obtain colours for the abnormal spectrum ‘E + A’ galaxies of Caldwell et al. (1993; 1996), many of which are to be found in the south-west corner of the cluster around a group of galaxies dynamically associated with NGC 4839 (Bailer 1984; Escalera, Sleazak & Mazure 1992; Colless & Dunn 1996). Significant advances in detector technology since the work of BLE92a allow us to use much larger CCDs with greater $U$-band sensitivity. In order to cover the required area, we took tiled images giving us continuous coverage of the cluster. In contrast, BLE92a targeted individual galaxies. Because of the continuous coverage, our sample of galaxies is more complete than that of BLE92a, including all of the GMP galaxies within our area of sky and with significantly higher precision than the GMP data.

### 2 OBSERVATIONS

The observing runs which provided data for use in this project are summarized in Table 1. $U$- and $V$-band observations were obtained in two successive years at the SAO 1.2-m telescope on Mt Hopkins, Arizona. The detector used was a thinned, back-side-illuminated, AR-coated 2048 x 2048 Loral CCD in 2 x 2 binning mode, giving us a 10-arcmin (192 $h^{-1}$ kpc) field of view with 0.63-arcsec pixels. The quantum efficiency of the CCD stays high and almost constant right across the $U$ band, giving effective filter responses which approximate the standard shape. The same setup was used for all observations to maintain a common photometric system for the whole data set. The average integration times used were 400 s in the $V$-band and 4 min in the $U$ band and the median $V$-band; seeing achieved throughout the run was 2.2 arcsec FWHM (see Appendix A).

The observations cover a continuous region encompassing the south-west group around NGC 4839, the central parts around NGC 4874 and 4889 and also a large amount of the north-east of the cluster (see Fig. 1). The observations were also designed to cover all of the Caldwell et al. (1993; 1996) abnormal spectra galaxies, however inclement weather meant that some were missed towards the extreme south-west and north-east. The presence of a seventh magnitude star just north of the centre was also avoided as

| Dates | Observer(s) | Usable Nights |
|-------|-------------|---------------|
| 20–21 March, 1996 | Caldwell | 1.5 |
| 11–14 April, 1996 | Caldwell & Terlevich | 4 |
| 9–11 May, 1996 | Caldwell | 4 |
| 1–5 April, 1997 | Caldwell & Terlevich | 1.5 |

Figure 1. The distribution of observed images across Coma. All observed galaxies with $V_{16} > 18$ are shown as a dot. The large dots show, from left to right, the positions of NGC 4889, 4874 and 4839 respectively. The dynamical centre of the Coma cluster is somewhere between NGC 4889 and 4874, while the dynamical centre of the substructure in the south-west corner is NGC 4839.
scattered light here makes data reduction difficult. To reduce any systematic differences in the photometry between parts of the cluster, observations of the central and south-west regions were interleaved during the observing runs.

During the night, immediately after dusk and before twilight, standard stars from Landolt (1992) were observed over a wide range in airmass. Care was taken to ensure the colours of the stars matched those of our galaxies, typically $0 \leq (U - V) \leq 2$. With our large field of view it was possible to observe many standards simultaneously. To have additional checks on the overall homogeneity of the final photometry, we used large overlaps of $\sim 1$ arcmin between the actual Coma cluster images, ensuring that the objects in this overlap region were observed in both images. We also interleaved snapshots (300 s and 100 s for $U$ and $V$ respectively) of the central parts of the cluster, thus using the galaxies there as ‘standard’ galaxies. This is particularly important for the $U$ band, as the spectral energy distribution of the standard stars is different from that of the early-type galaxies.

3 DATA REDUCTION

The images were reduced using the standard methods in the IRAF package. The CCD used had a number of cosmetic defects, so in the subsequent reduction procedures we mark objects within 5 pixels of a defect as suspect.

3.1 Galaxy identification, photometry and astrometry

Lists of candidate objects were produced using the Sextractor 1.2b10 program (Bertin & Arnouts 1996) from the V-band images. Sextractor was also used to differentiate the extended sources from the stellar sources in the images. All objects within 15 pixels (9.45 arcsec) of a CCD edge were rejected. In order to avoid problems matching galaxies on the $U$-band frames, galaxies fainter than $V = 20$ were not carried forward for further analysis. This gives rise to a sharp cut-off in the magnitude distribution of our catalogue.

The list of positions for the galaxies in each image was used by the IRAF phot package to generate fixed aperture magnitudes in 8.8–30 arcsec diameter apertures. The sky level was measured in annuli with inner radii of 50 arcsec. For the fixed photometric apertures with radii less than 20 arcsec, we also measured the sky level at 25 arcsec, and we give both values in the final data table. The magnitudes measured were then corrected for the varying seeing between the frames as described in Appendix B. Astrometry for the frames was calculated using the HST guide star catalogue as a source of reference stars. The rms scatter in our astrometry is approximately 1 arcsec. We have not corrected the data for geometric distortions in our detector, as they introduce photometric errors of less than 1 per cent between galaxies at the centre and galaxies at the edge of the detector.

Observations of standard stars left residuals at the 0.03 mag level in the $U$ band, once a drift in the zero-point had been corrected. The drift (in all cases less than 0.2 mag per night) is also evident when examining the overlap regions between images and the repeated observations of ‘standard’ Coma fields. As these offsets were so readily measurable, we used them to improve the zero-points of the individual images in order to ensure that the whole cluster is on a consistent photometric system for all the observations. The method we used to generate this system is very similar to that used by Maddox, Efstathiou & Sutherland (1990), and is described in detail.

Table 2. This table lists the rms errors in our photometry for all of our apertures. It was generated using the repeated observations of both stars and galaxies, mostly in the regions of overlap between images. Fig. 2 shows how the rms errors vary with magnitude for the 8.8-, 13- and 20.2-arcsec diameter apertures. The rms internal scatter quoted by BLE92a for their 13-arcsec photometry, which only reaches a magnitude of $V_\odot = 16.5$, is 0.025 and 0.015 mag for $U$ and $V$ respectively.

| Aperture diameter (arcsec) | $U$ RMS scatter | $V$ RMS scatter | $(U - V)$ RMS scatter |
|---------------------------|-----------------|-----------------|-----------------------|
| 8.8                       | 0.02303         | 0.01619         | 0.02829               |
| 12.6                      | 0.02124         | 0.01364         | 0.02532               |
| 13                        | 0.02258         | 0.01352         | 0.0264                |
| 16                        | 0.0277          | 0.01478         | 0.03145               |
| 20.2                      | 0.03036         | 0.01727         | 0.035                 |
| 25.2                      | 0.0359          | 0.02076         | 0.04154               |
| 26                        | 0.04805         | 0.02706         | 0.05528               |

Figure 2. In the three panels, we plot the observational errors as a function of $V_\odot$ magnitude, from repeated measurements of galaxies in overlapping regions. The $(U - V)$ colour errors are shown in panel (a). Panels (b) and (c) show the errors for the $U$ and $V$ photometry. The three lines show a running biweight scatter (see Section 3.2) measured through circular apertures of 8.8, 13 and 20.2 arcsec diameter. A bin size of 60 observations was used in calculating the running mean, reducing to six at the bright extreme of the plot. Table 2 lists the mean rms scatter down to $V_\odot = 17$ mag for all of our apertures. The increasing levels of scatter with aperture size is entirely consistent with the increasing contribution of the sky to the noise with aperture radius.
in Appendix B. The absolute calibration of this system was set to agree with the published $U$- and $V$-band photometry of BLE92a.

### 3.2 Quantifying the photometric errors

We can use the large number of multiply observed objects (mostly from the large overlaps between images) to constrain our observational errors. Table 2 lists the mean rms scatter for all objects with $V_{13} < 17$ mag in all of our apertures, and Fig. 2 shows how the scatter in the observations of multiply observed objects increases with magnitude for the 8.8-, 13- and 20.2-arcsec diameter apertures. The scatter was computed using a running biweight scatter indicator (Beers, Flynn & Gebhardt 1990). We use the biweight scatter indicator throughout this paper owing to its robustness and resistance in the case of non-Gaussian distributions. Additionally, we perform all regression analysis by minimising both the biweight scatter and location (mean) of the residuals to the fit. This provides an efficient way of performing fits to data which in the case of the CMR residuals have a non-Gaussian distribution, and often a large tail.

The levels of scatter in our 13-arcsec apertures are the same as those quoted by BLE92a (0.025 and 0.015 mag for $U$ and $V$ respectively), and they stay constantly low down to $V_{13} \sim 17$ mag.

It should be noted that although Table 2 shows that the 12.6-arcsec aperture has the lowest rms scatter in both $U$ and $V$ bands, Fig. 2 shows that fainter than $V_{13} \sim 17$, it is quickly overtaken by the 8.8-arcsec aperture, which is better suited to the smaller sizes of the fainter galaxies.

As an independent check of our calibration, we have compared our photometry directly with that of BLE92a (see Fig. 3). The scatter between our colours and theirs is 0.034 mag, while the scatter between our photometry and theirs is 0.022 and 0.032 mag for the $V$ and $U$ bands respectively. This is almost exactly what we expect simply by adding the rms internal scatters of our data and theirs in quadrature. Equally important, given the method used to obtain a uniform photometric system for our data, is the fact that the mean colour difference does not vary as a function of distance from NGC 4874 (Fig. 3a).

### 3.3 The photometric diameter $D_V$

The photometric diameter parameter, $D_V$, is equivalent to the $D_n$ parameter used by Dressler et al. (1987), but is based on $V$-band photometry. We use the definition of $D_V$ given in Lucey et al. (1991), namely that $D_V$ is the photometric diameter (in arcsec) that encloses an area of average surface brightness of 19.80 mag arcsec$^{-2}$. Like those of Lucey et al., our $D_V$ values include a $(1+z)^{4}$ cosmological correction. As the $D_V$ are (mostly) calculated using interpolation, they are very accurate. In order to measure $D_V < 8.8$ arcsec (our smallest photometric aperture), we also need to perform some extrapolation, and we use $D_{8.8}$ only down to 4 arcsec to keep this to a stable minimum. The function we use for both interpolating and extrapolating is a simple $R^{1/4}$ profile, which is fitted to the seeing-corrected aperture magnitudes. In addition to rejecting $D_{8.8}$ smaller than 4 arcsec, we also do not attempt to calculate $D_{8.8}$ for galaxies where the function fit was poor, or where extrapolation of the data to larger radius is necessary. Comparison with the independent $D_V$ values given by

![Figure 3](https://example.com/figure3.png)

**Figure 3.** The four panels show the behaviour of the residuals between our 13-arcsec diameter aperture photometry and that of BLE92a. In all cases, the residuals are calculated by subtracting the BLE92a data from ours; for example, a negative $\Delta(U-V)$ indicates that our colour for a galaxy is bluer than the BLE92a colour. The solid and dashed lines show running biweight location and scatter indicators. Panel (a) shows the difference between the $(U-V)_{13}$ colours obtained in this paper and those of BLE92a, as a function of distance from NGC 4874. Panel (b) shows how our colours compare with the BLE92a colours as a function of luminosity. Panels (c) and (d) show how our $U$- and $V$-band magnitudes compared with those of BLE92a.
Lucey et al. (1991) for the same galaxies shows the uncertainty to be better than 0.007 dex.

3.4 Catalogue

An extract of the photometric catalogues are presented in Table 3. The full version, which is in the on-line version of the article in Synergy, available electronically, contains $U$- and $V$-band photometry for all the apertures listed in Table 2.

4 ANALYSIS TECHNIQUE

The first step in our analysis is to establish a criterion for cluster membership. Redshifts are available for all galaxies brighter than $V_{13} = 15.7$. For the fainter galaxies, 96 per cent of galaxies with $V_{13} < 16$, 89 per cent of $V_{13} < 17$ galaxies, and 71 per cent of $V_{13} < 18$ galaxies have redshifts. The velocities were obtained from the NASA/IPAC Extragalactic Database (NED). Most of the velocities can be attributed to Colless & Dunn (1996). Galaxies with recessional velocities between 4000 and 10 000 km s$^{-1}$ were taken as confirmed members of the Coma cluster. From this sample we then removed galaxies with ‘bad’ photometry. Galaxies were deemed bad, and thus rejected if

(i) emission from a nearby object entered the 13-arcsec diameter aperture,
(ii) one of the CCD bad columns passes through the galaxy,
(iii) a cosmic ray was removed from part of the galaxy in the $V$ image (cosmic rays in the $U$ images were less of a problem owing to the image being composed of multiple exposures).

The morphological classification of galaxies is taken from Andreon et al. (1996) and Andreon, Davoust & Poulain (1997) which gives morphologies for all galaxies. Andreon et al. (1996) and Andreon, Davoust & Poulain (1997) to the image being composed of multiple exposures).

Table 3. A small extract of the photometric catalogue. Owing to its size, the full catalogue is only available electronically, and can be found in the electronic version of the article in Synergy.

| GMP RA J2000 | DEC J2000 | log$_{10}(D_V)$ | $V_{13}$ | $U_{13}$ | $V_{16}$ | $U_{16}$ |
|-------------|-----------|----------------|---------|---------|---------|---------|
| 1807 13:01:50.2 | 27:53:36.2 | 0.943 | 15.322 | 16.727 | 15.161 | 16.570 |
| 1853 13:01:47.0 | 28:05:41.5 | 1.055 | 14.866 | 16.317 | 14.704 | 16.167 |
| 1885 13:01:44.1 | 28:12:51.4 | 0.628 | 16.747 | 17.957 | 16.630 | 17.845 |
| 2000 13:01:31.8 | 27:50:50.9 | 1.112 | 14.636 | 16.105 | 14.456 | 15.924 |
| 2048 13:01:27.2 | 27:59:56.8 | 0.821 | 16.039 | 17.435 | 15.938 | 17.338 |
| 2059 13:01:26.2 | 27:53:09.9 | 1.070 | 14.774 | 16.288 | 14.565 | 16.099 |

Table 4. The morphological classes used in this paper, combined into the broad categories of late-type, S0 and early-type. Throughout the paper, no distinction shall be made between discy, boxy and undefined ellipticals (diE, boE, unE). They shall simply be referred to as ellipticals.

| Category | Morphological type | Number of galaxies |
|----------|-------------------|--------------------|
| late-type | Sp, Sb, Sa, SA0/a, SB0/a, SAB0/a, SAB0p | 32 |
| S0 type | SA0, SAB0, SB0, diE/SA0, diE/SAB0, unE/SA0 | 71 |
| early type | Epec, boE, diE, unE | 26 |
| No morph | | 148 |

The morphological classifications are taken from Andreon et al. (1996) and Andreon, Davoust & Poulain (1997) which gives morphologies for all $V_{13} < 18$ galaxies have morphological information. The elliptical and late-types are uniformly distributed throughout this magnitude range, however the proportion of S0 galaxies relative to the total number has a sharp peak at $V_{13} = 16$ mag, where 68 per cent of the galaxies are S0s. The different symbols in Figs 4, 6, 7 and 10 correspond to the broad morphological type of the galaxy. The actual morphological types used in these broad classifications are shown in Table 4, together with the frequency of each type.

Regression analysis of the colour–magnitude relation was performed using the biweight estimator. Our aim is to distinguish the ridge-line of the CMR, and we do not want to be unduly influenced by the exact position of blue outliers. The biweight is a robust, resistant and efficient location and scale indicator (as is apparent from Table 6 later). We compared the results of the biweight technique to that of using a biweight technique in conjunction with a $3\sigma$ clipping of the data set, however the difference was within the $1\sigma$ error estimate of the unclipped method.

The errors in the best-fitting relation were calculated by bootstrap resampling of the data. The observational uncertainties in the colour as a function of galaxy luminosity $[O,(L)]$ are well known for this data set (see Fig. 2), so in order to make an estimate of how much of the measured scatter in the CMR is the result of observational errors, we defined a mean observational colour scatter thus:

$$\sigma_c = \frac{\sum_{i=1}^{N} O_c(L_i)}{N},$$

where the $L_i$ are the luminosities of the galaxies in the data set, and $N$ is the number of galaxies. Using this value for the observational errors, a value for the intrinsic scatter in the CMR can be calculated:

$$I = \sqrt{\sigma^2 - \sigma_c^2},$$

where $\sigma$ is the observed scatter of the CMR.

Owing to the fact that we are measuring the CMR scatter using a limited sampling of the underlying distribution, some uncertainty is introduced. This can make the measured scatter ($\sigma$) smaller than...
the observational errors \( (\sigma) \). In such a cases, we show the intrinsic scatter as zero, but we can still show an upper limit to the intrinsic scatter from the bootstrap limits.

5 ENVIRONMENTAL AND MORPHOLOGICAL VARIATIONS

Fig. 4 shows the colour–magnitude relation for every extended object in the photometric catalogue. Many of the objects shown will not be members of the Coma cluster, yet despite this, the CMR is clearly visible down to \( V_{25.2} = 19 \) mag. We are interested in measuring changes in the CMR, such as in its scatter or slope, in different parts of the cluster, in different subsets of galaxy morphology, and for different luminosities. We therefore concentrate solely on those galaxies identified as members of the cluster from their recessional velocity. Using the recessional velocity avoids the need for statistical background subtraction. We have used the properties of the galaxies, such as morphology, luminosity and position, to define 14 subsets of these 275 member galaxies. The subsets are defined in Table 5.

Fig. 6 shows the CMR for all confirmed cluster members (data set 1). The CMR is made up of galaxies of differing morphological types, and from every part of the cluster, yet it extends for almost five magnitudes without deviating from a straight line. Fig. 4 shows that the CMR actually extends fainter than this in our data, but we have no redshifts for these faint galaxies. Secker, Harris & Plummer (1997) have shown that the \((B, B – R)\) CMR in dwarf ellipticals in Coma actually continues down to at least \( B \sim 21.5 \). Another important aspect of Fig. 5 is in the direction of scatter. There is almost no scatter redward of the CMR ridge line, even at the faint end where the observational errors are greatest, there is however significant blueward scatter, most of which is caused by the late-type population (Fig. 6).

In the following sections, we investigate the properties of the

![Figure 4](image_url)

**Figure 4.** The \((U,V)\) colour–magnitude relation for all objects detected for which SExtractor gives a CLASS_STAR \( \geq 0.2 \). The \( V \) are taken from the 25.2-arcsec diameter aperture, as it is more complete than the 26-arcsec aperture. In order to increase the signal-to-noise ratio in the colour term, the \((U – V)\) are taken from the smallest aperture, the 8.8-arcsec diameter aperture. The symbols represent the morphological types of the galaxies from Andreon et al. (1996, 1997). The small open symbols have no morphological information. The line is a biweight fit (see Section 3.2) to the data, and follows the ridge line of the CMR, which can be seen to extend down to \( m_{V_{25.2}} \sim 19.5 \).

| Data set | Description |
|----------|-------------|
| 1        | Confirmed members: After rejecting the ‘bad’ galaxies (see Section 4), we classify all galaxies from 4000 km s\(^{-1}\) to 10 000 km s\(^{-1}\) as members of the cluster. The velocities were obtained from the NASA/IPAC Extragalactic Database (NED). Most of the velocities can be attributed to Colless & Dunn (1996). |
| 2        | Confirmed members with \( V_{17} < 17 \): a subsample of confirmed members (data set 1) with the faint tail cut off at the point where the measurement errors in the colours starts to increase (see Fig. 2). |
| 3        | All with morphology: all member galaxies with a morphology from Andreon et al. (1996; 1997) (see Table 4). |
| 4        | E&S0 morphology: all member galaxies with an elliptical or S0 morphological type (see Table 4). |
| 5        | S0 morphology: all member galaxies with S0 morphology (see Table 4). |
| 6        | Elliptical morphology: all member galaxies with elliptical morphology (see Table 4). |
| 7        | Late-type morphologies: all member galaxies with late-type (spiral and irregular) morphology (see Table 4). |
| 8        | E&S0 centre: early-type galaxies closer to NGC 4874 than 4839. |
| 9        | E&S0 SW: early-type galaxies closer to NGC 4839 than 4874. |
| 10       | E&S0 inner: early-type galaxies within 15 arcmin of either NGC 4874 or 4839. |
| 11       | E&S0 outer: early-type galaxies further than 15 arcmin from both NGC 4874 and 4839. |
| 12       | E&S0 bright: the bright half of data set 4. |
| 13       | E&S0 faint: the faint half of data set 4. |
| 14       | Members bright: the bright half of data set 1. |
| 15       | Members faint: the faint half of data set 1. |
CMR in each of the data sets (see Fig. 7). We use the techniques described in Section 4 to ascertain the scatter about the main ridge line of the CMR, as well as the scatter in the total sample. Throughout the rest of the paper we use the 13-arcsec diameter aperture magnitudes and colours (resulting from their low photometric errors) or colours measured within the $D_V$ diameter (in order to define colours within an aperture that scales in galaxy size).

In order to verify whether the findings of the following sections can be attributed to variations in the CMR, rather than selection biases in the sample, we have investigated luminosity and morphological segregation within the samples. Using a Kolmogorov-Smirnov (K-S) test, we find that the late-type and S0 samples have statistically indistinguishable luminosity distributions, but that the early-type sample is on average 1 mag brighter. We also find no correlation between luminosity and angular distance from NGC 4874, nor any significant difference between the luminosity distribution of the E&S0 inner and E&S0 outer data sets (see Fig. 11 later).

### 5.1 Morphological dependence of the CMR

In this section we examine variations in the CMR of galaxies of different morphological types. We use the broad morphological types defined in Table 4. The dividing line between the various...
types is somewhat arbitrary, and we err towards the later morphological types, i.e. we classify a galaxy of type E/SA0 as S0, and one of type SA0/a as late type. We investigate the scatter of the main ridge line of the CMR and the amount of blueward scattering separately. Initially we concentrate on the ridge line. Looking at the morphologically segregated data sets (3,4,5,6 and 7) in Table 6, they all have levels of intrinsic scatter indistinguishable within the measurement errors (\(\sim 0.05\) mag), except for the elliptical galaxies (data set 6) which has significantly lower levels of scatter (0.036 mag).

The only exception is the late type galaxy data set (7): however, even this data set includes many objects that lie on the CMR ridge line. Six of the late-type galaxies are very blue compared with the CMR ridge-line. The presence of these blue galaxies is not surprising. Fig. 6 shows snapshots of the V-band images of these galaxies, which even with our poor spatial resolution can be made out to be very obviously late-type. The only S0 galaxy amongst these, NGC 4853, was identified by Caldwell et al. (1993) as a post-starburst galaxy, and is included here with the late types because of peculiar asymmetries in its light profile. Perhaps the most surprising aspect of these blue galaxies is not their presence, but the fact that there are only seven of them, out of the 32 late-types in our sample. If we apply a 3\(\sigma\) clipping to this data set, we obtain a CMR ridge-line that is indistinguishable from the S0 types. Fig. 8 shows snapshots of the V-band images of these ‘red’ late-types. All of the galaxies in Fig. 8 are within 3\(\sigma\) of the best-fitting CMR for data set 6. Some of the galaxies are borderline SA0/a, however they are the minority, and cannot in themselves

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**Figure 7.** This figure (continued on pages 9 and 10) shows the CMRs for each of the data sets defined in the main text. The symbols are the same as those used in Fig. 6. Dashed lines represent the 3\(\sigma\) scatter of the galaxies about the best-fitting line for the full data set.
explain the low scatter. All of the ‘red’ late type galaxies appear to have fainter discs compared with the ‘blue’ ones in Fig. 6, however statistical tests show no correlation between the $D_e$ of a galaxy and its CMR residual $[U - V] - (mV + c)$, where $m$ and $c$ are the best-fitting slope and intercept for the CMR). There is however a strong correlation between CMR residual and colour gradient, and whereas the majority of galaxies have colour gradients in the sense that they become bluer with increasing radius, the blue galaxies have a colour gradient in the opposite direction (they become redder with increasing radius). Thus, although the ‘blue’ galaxies are not necessarily more compact (cf. Moss & Whittle 2000), the ‘blue’ light is more concentrated than the CMR galaxies. We conclude that the ‘red’ late-types are ‘anaemic’ (e.g. Van Den Bergh 1991), i.e. that they have lost their H I gas through interactions with the intracluster medium.

The only other data set with large numbers of ‘blue’ galaxies is data set 1 (all cluster members). These blue galaxies are, in addition to the late-type galaxies noted above, morphologically untyped. Fig. 9 shows V-band images of all of the galaxies that deviate from the CMR ridge line for data set 1 by more than 5σ. It is immediately obvious that they are predominantly of late type (e.g. GMP 4570), although in some cases it is difficult to tell (e.g. GMP 3848).

### 5.2 Luminosity dependence of CMR

The last set of data sets are the ones where we segregated the galaxies according to their luminosity (12, 13, 14 and 15). We have separated the early-type galaxies data set (4) into two halves of equal numbers, with the bright half of the galaxies in data set 12, and the faint half in data set 13. Because data set 1 spans a much larger range in luminosity than do the data sets with only morphologically typed galaxies, we split that into two halves too (data sets 14 and 15).

The bright and faint early-type galaxies have indistinguishable slope and intercept, however the faint sample has greater intrinsic scatter. This could be related to the increasing numbers of blue fainter magnitudes. As these galaxies start to become apparent after $V_{13} > 16$, they are mainly untyped in our study.

The two halves of the complete members data set have differing slopes. It is possible that this could simply be the result of aperture effects, but as we do not have reliable $D_e$ measurements for the faint half of the data set, we cannot check this directly (see below). Although the observed scatter in the faint sample is larger than that of the bright sample, it has twice the mean observational error ($\bar{O}_s$). Once the difference in the observational errors is taken into account, both the bright and the faint samples have similar values for the intrinsic scatter.

A common concern over the interpretation of the CMR as a constraint on galaxy formation (e.g. Bower et al. 1998; Kauffmann & Charlot 1998) is with the role played by colour gradients within galaxies. Is it possible, for example, that the slope of the CMR is caused by the metric aperture measuring a larger fraction of the total light in small galaxies than in larger ones? To address this effect, we have measured the colour within the $D_e$ diameter of each galaxy. This gives a measure of the colour of each galaxy that scales with the properties of the galaxy. This approach is preferable to measuring the colour within a fixed fraction of the total light because $D_e$ is (usually) a radius at which the colour can be accurately defined and does not require extrapolation. In contrast, measurements based on the effective radius, $R_e$, require extrapolation of the radial profile, and often require the colour to be measured at a radius where the signal-to-noise ratio is low.

The colour–magnitude relation measured within $D_e$ is shown in Fig. 10. The slope of this relation is weaker than that measured within the fixed diameter, though it is still clearly evident. Table 7 shows the slope of the relation for the whole sample and for the early-type data sets. Bower at al. (1998) estimated that the approximately one third of the CMR slope resulted from aperture effects. Comparison with Table 6 shows that the reduction in the slope is consistent with this. This data set therefore confirms that the existence of the CMR slope is not an artefact arising from radial colour gradients in the galaxies. Any model of cluster galaxy formation must therefore be able to explain simultaneously the small scatter of the relation and its slope. Importantly, this constrains the factor by which the masses of cluster galaxies can grow through random collisions between objects of different colours (see Bower et al. 1998 for a fuller discussion).

### 5.3 Environmental dependence of the CMR

In order to investigate the environmental dependence of the CMR, we have defined subsamples of the cluster members according to their position in the sky. Datasets 10 and 11 contain galaxies that

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**Table 6. Results of regression analysis on the 13-arcsec diameter aperture magnitudes.** The errors quoted for the slope, intercept and observed scatter are 1σ bootstrap errors. The mean observational scatter ($\bar{O}_s$) is used in calculating the intrinsic scatter (see main text).

| Data set | Number | Slope | Intercept | Observed scatter (σ) | $\bar{O}_s$ | Intrinsic scatter (σ) |
|----------|--------|-------|-----------|----------------------|------------|----------------------|
| (1) All members | 275 | $-0.128 \pm 0.001$ | $3.37 \pm 0.1$ | 0.0744 | 0.0658 | 0.0835 | 0.029 | 0.069 | 0.059 | 0.078 |
| (2) $V_{13} < 17$ | 175 | $-0.139 \pm 0.003$ | $3.55 \pm 0.2$ | 0.0674 | 0.0585 | 0.0751 | 0.025 | 0.063 | 0.053 | 0.071 |
| (3) All with morph | 129 | $-0.142 \pm 0.003$ | $3.59 \pm 0.2$ | 0.0599 | 0.0506 | 0.0675 | 0.024 | 0.055 | 0.044 | 0.063 |
| (4) E&S0 morph | 97 | $-0.137 \pm 0.003$ | $3.5 \pm 0.3$ | 0.0552 | 0.0471 | 0.0638 | 0.024 | 0.05 | 0.041 | 0.059 |
| (5) S0 morph | 71 | $-0.125 \pm 0.007$ | $3.32 \pm 0.4$ | 0.0583 | 0.0453 | 0.0683 | 0.023 | 0.053 | 0.039 | 0.064 |
| (6) E morph | 26 | $-0.164 \pm 0.006$ | $3.91 \pm 0.5$ | 0.0438 | 0.0298 | 0.0504 | 0.025 | 0.036 | 0.016 | 0.044 |
| (7) S0p morph | 32 | $-0.166 \pm 0.001$ | $3.96 \pm 0.7$ | 0.0778 | 0.0458 | 0.122 | 0.025 | 0.074 | 0.038 | 0.12 |
| (8) E&S0 center | 78 | $-0.132 \pm 0.004$ | $3.43 \pm 0.3$ | 0.0517 | 0.0435 | 0.0584 | 0.024 | 0.046 | 0.036 | 0.053 |
| (9) E&S0 SW | 19 | $-0.15 \pm 0.009$ | $3.71 \pm 2$ | 0.0755 | 0.0485 | 0.119 | 0.024 | 0.069 | 0 | 0.12 |
| (10) E&S0 inner | 55 | $-0.149 \pm 0.005$ | $3.71 \pm 0.3$ | 0.0542 | 0.0432 | 0.0636 | 0.024 | 0.049 | 0.036 | 0.059 |
| (11) E&S0 outer | 42 | $-0.124 \pm 0.003$ | $3.28 \pm 0.3$ | 0.0412 | 0.0282 | 0.0532 | 0.024 | 0.033 | 0.014 | 0.047 |
| (12) E&S0 bright | 49 | $-0.135 \pm 0.007$ | $3.49 \pm 0.6$ | 0.0461 | 0.0366 | 0.0528 | 0.025 | 0.039 | 0.027 | 0.046 |
| (13) E&S0 faint | 49 | $-0.135 \pm 0.009$ | $3.48 \pm 0.7$ | 0.065 | 0.049 | 0.081 | 0.023 | 0.061 | 0.043 | 0.078 |
| (14) members bright | 140 | $-0.141 \pm 0.005$ | $3.57 \pm 0.3$ | 0.0657 | 0.0557 | 0.075 | 0.024 | 0.061 | 0.05 | 0.071 |
| (15) members faint | 136 | $-0.123 \pm 0.004$ | $3.29 \pm 0.3$ | 0.086 | 0.069 | 0.107 | 0.054 | 0.067 | 0.043 | 0.093 |
are closer to and further away, respectively, than 15 arcmin from NGC 4874, while data sets 8 and 9 contain galaxies that are nearer to NGC 4874 or 4839 respectively (see table 5). We have restricted ourselves to early-type galaxies (ellipticals and S0s) to avoid, as much as possible, any bias arising from the morphology–density relation within the cluster, although Section 5.1 showed there to be little variation in the properties of the CMR ridge line between the different morphological types.

5.3.1 Significance of the environmental change

The low number of galaxies in the south-west sample makes it very difficult to measure its scatter. The measured scatter in the full data set is $0.07 \pm 0.06$, so we instead concentrate on the scatter in the central and outer samples (data sets 10 and 11). Here we see something unexpected. The outer data set has less scatter (both observed and intrinsic) than the inner data set. Although it is only

![Figure 8](https://academic.oup.com/mnras/article-abstract/326/4/1547/1068551)
marginally significant, it is just the opposite of what we would expect. Fig. 7 (panels 10 and 11) show the CMRs for these data sets. From them we can see that they both have similar numbers of elliptical and S0 galaxies. There are too few elliptical galaxies in each data set to be able to measure the scatter reliably for just the ellipticals in each one, but it is possible for the S0 galaxies. This shows the effect to be unrelated to morphology, with the 41 inner S0 galaxies having an observed scatter of 0.0597 ± 0.02 and the 30 outer galaxies having an observed scatter of 0.0377 ± 0.01. Fig. 11 shows the luminosity distribution for both data sets, which are statistically indistinguishable using both a K–S test and a Student’s t test (see Section 5).

Another approach to analysing the environmental dependence of the CMR is to analyse the data by averaging over radial bins. Because there are insufficient data within a single bin, we use a running biweight to analyse the radial dependence. We have excluded the galaxies around the NGC 4839 group in order to ensure that the radial sample also corresponds to a gradient in density. We ordered the galaxies according to their projected distance from NGC 4874, and calculated running biweight location and scale indicators for the residuals from the CMR ridge line for three data sets: the full data set (1), the early-type galaxy data set (4) and a data set of just the S0 galaxies (5) (see Fig. 12). In all three panels, the \((U - V)\) residuals are calculated using the best-fitting CMR to the whole data from Table 6, such that negative values of \(\Delta(U - V)\) imply a galaxy is positioned to the blue side of the CMR ridge line. The first panel, using all the cluster members, shows no increase in scatter with radius, however this could be affected by greater numbers of late-type galaxies in the outer parts of the cluster. We therefore also show the colour residuals of the early-type galaxy data sets (4 and 5). Again the scatter remains almost constant, with a very small downward gradient. All three

Figure 9. V-band images of all of the galaxies that lie more than 5σ blueward of the CMR for the members data set (data set 1). In addition to the galaxies identified as starburst or post-starburst in Fig. 6 (also present in this figure) GMP 4255 (D44) is identified as a post-starburst and GMP 4579 (D45) is identified as a starburst by Caldwell et al. (1993). Text and symbols in the plot are as for Fig. 8.
panels show that the residuals from the CMR seem to become systematically bluer towards the edges of the cluster. It seems unlikely that this could be an age effect without also incurring an increase in the scatter of the CMR, which leaves two possibilities. First, it could simply be a radial drift in our photometric zero points. Although we checked for this against the data of BLE92a (see Fig. 3), it only extends out to a radius of 15 arcmin, which is also where this effect begins to be noticeable. We can however make a quick estimate of the expected drift in our photometry. We have on average 20 objects in the overlap regions, with rms photometric errors of 0.026 mag (see Table 2), which gives an rms colour error between images of $0.026/\sqrt{20}$. Now to get to a radius of 30 arcmin, we need to traverse at least three image boundaries, so the error accumulated is $\sqrt{3} \times 0.026/\sqrt{20} = 0.01$ mag. However, the radius can be calculated in many different directions, and the photometric zero-point for each image was indeed calculated in an iterative way such that any errors would dissipate in a two-dimensional manner, so the value of 0.01 mag (much smaller than the value of the colour gradient) can be regarded as an upper limit.

5.3.2 Dust

A possible explanation for the gradient could be that the galaxies in the centre of the cluster are being reddened because of intracluster dust. The upper limit on the reddening through dust in the core of Coma, in comparison with the field is $E(U-V) \leq 0.08$ mag. Ferguson (1993) could account for this amount of reddening; it could also add an extra source of scatter to the central parts of the cluster not present in the outer parts. Galaxies behind the cluster would appear both fainter and redder than an identical galaxy in front of the cluster. This would tend to increase the scatter in the CMR, but only in the central parts.

We can estimate the contribution of this dust to the scatter in the core of the cluster as follows. We define $r$ as the distance from the core along the line of sight. If $r_c$ is the virial radius of the cluster, then $r = -r_c$ is the front of the cluster, and $r = r_c$ is the rear of the cluster. We also assume that the galaxies and dust have the same isothermal density distribution, out to the virial radius

$$
\delta(r) = \begin{cases} 
1 & \text{if } |r| \leq r_c, \\
1 + \left(\frac{r}{r_c}\right)^2 & \text{if } |r| > r_c,
\end{cases}
$$

Figure 10. The colour–magnitude relation measured with in the photometric diameter $D_p$. In this plot, the area within which the photometry is measured varies according to galaxy size. $V_{D_p}$ is the V-band magnitude within the $D_p$ diameter; $(U - V)_{D_p}$ is the colour within this diameter.

Figure 11. The luminosity distributions for the inner and outer E + S0 data sets (10 and 11). The lined histogram represents the luminosity distribution for the outer data set (11) and the outlined histogram represents the luminosity distribution of the inner data set (10). Both distributions are statistically indistinguishable using both a K–S test and a Student’s t test (see Section 5).

Table 7. Results of regression analysis on the aperture magnitudes measured within $D_p$ diameter apertures. For a complete description of each data set see the main text. The errors quoted for the slope are 1σ bootstrap errors.

| Dataset | Number | Slope |
|---------|--------|-------|
| (3) All with morph | 111 | $-0.0754 \pm 0.005$ |
| (4) E&S0 morph | 86 | $-0.0819 \pm 0.004$ |
| (5) S0 morph | 63 | $-0.082 \pm 0.006$ |
| (6) E morph | 23 | $-0.0911 \pm 0.003$ |
stellar population of age 10 Gyr from Bower et al. (1998), we find

\[ \text{s} \quad \text{the other possibility for the bluing towards the edges is a} \]

5.3.3 Age

than the number quoted by Ferguson (1993).

elliptical galaxies, we can say that either the dust is not distributed

It should be remembered that this is a very rough model for the
distribution of dust and galaxies in the cluster. The distribution of
dust especially is very poorly known. Clearly from the fact that the
levels of scatter in this model is greater that that measured for
elliptical galaxies, we can say that either the dust is not distributed
as an isothermal sphere, or the limit for dust in the core is lower
than the number quoted by Ferguson (1993).

\[ \text{where} \quad r_c \quad \text{is the core radius. The amount by which a galaxy at} \quad r \quad \text{is}
\]

\[ R(r) = \alpha r_c \left[ \arctan \left( \frac{r}{r_c} \right) + \arctan \left( \frac{r_c}{r} \right) \right]. \]

\( \alpha \) is a constant, which we chose in order to satisfy our boundary
condition, so that we get the required reddening in the centre of the
cluster, i.e. \( R(0) = 0.08 \) mag (Ferguson 1993). We then calculate
the mean and standard deviation of \( R(r) \):

\[ \bar{R} = \alpha r_c \arctan \left( \frac{r_c}{r} \right) = 0.08 \text{ mag} \]

and

\[ R^2 = 4 \left( \frac{4}{3} \alpha r_c \arctan \left( \frac{r_c}{r} \right) \right)^2 = \frac{4 \times 0.08}{3} \text{ mag}^2. \]

It should be noted that because of our choice of distributions for the
dust and galaxies, both \( r_c \) and \( r_r \) have cancelled out of the
calculations. The standard deviation of \( R(r) \) is then

\[ SD(R) = \sqrt{R^2 - R^2} = 0.046 \text{ mag}. \]

It should be remembered that this is a very rough model for the
distribution of dust and galaxies in the cluster. The distribution of
dust especially is very poorly known. Clearly from the fact that the
levels of scatter in this model is greater that that measured for
elliptical galaxies, we can say that either the dust is not distributed
as an isothermal sphere, or the limit for dust in the core is lower
than the number quoted by Ferguson (1993).

5.3.3 Age

The other possibility for the bluing towards the edges is a
difference in mean galactic age. Using the models for a single burst
stellar population of age 10 Gyr from Bower et al. (1998), we find
that \( d(U - V)/dt \sim 0.03 \) mag Gyr\(^{-1} \). This would make the outer
galaxies approximately 2 Gyr younger than the central galaxies.
Assuming lower ages for the galactic population would make the
difference in age between the inside and the outside smaller, i.e. if
the galaxies are only 5 Gyr old, the difference in age between the
inner and outer galaxies is only 1 Gyr. Abraham et al. (1996) find a
\( (g-r) \) colour gradient with projected radius in the \( z = 0.23 \) cluster
Abell 2390 of \( m = -0.08 \) mag log\(_{10}(r_p) - 1 \), which they attribute to
an age trend. To compare the Coma colour gradient with that of
A2390, we used template early type galaxy spectra to K-correct
the Coma colours to the redshift of Abell 2390, and to convert
them from \( U - V \) to \( g - r \). We find that the gradient shown in
Fig. 12 for early-type galaxies is transformed into \( m = -0.08 \) mag log\(_{10}(r_p) - 1 \), a third of that measured in A2390.

A similar argument to the one above for the increased scatter in
the core arising from dust also applies in this case. When we look at
the core, we also include galaxies in the foreground and
background that are not in the cluster core, so are bluer than the
core galaxies. This effect is not as large as the dust effect, however,
because the galaxies behind the cluster are just as blue as galaxies
in front of it, so the effect is roughly half that expected from the
dust model.

6 CONCLUSIONS

We have placed new limits on the levels of scatter in the \((U, V)\)
CMR of the Coma cluster. The cluster members were split into
groups depending on their morphotype, luminosity or position on
the sky, and the CMR was studied in each of them. We found the
properties of the ridge line to be surprisingly consistent between all
of these groups. We have also calculated upper and lower limits for
the intrinsic scatter in each galaxy sample, taking into account the
low-number statistics that we are dealing with for some of them.
The results are presented in Tables 6 and 7.

We find no variation in the slope of the CMR ridge line between
elliptical and S0 morphological types. The late-type galaxies in the
cluster have a marginally steeper slope. This could be connected to
the increased blue scatter we find towards the faint end of the
CMR. In the galaxies for which we have morphological types, all
of these very blue galaxies are late-types. All of the ‘blue’ galaxies,
even where we do not have morphological data, have colour
gradients in the opposite direction to the normal CMR galaxies,
i.e. the blue galaxies get redder with increasing galactic radius.
This could be the result of star formation constrained to the
galaxy core and agrees with the findings of Moss & Whittle
(2000), who show that a disturbed cluster galaxy morphology is
a strong predictor of compact H\alpha emission. Fig. 9 shows that
even with our poor spatial resolution, which tends to make
galaxies appear to be of an earlier type, many of the unclassified
blue galaxies are also of late type.

Although the bluest galaxies tend to have late-type morphology,
there are many galaxies with late-type morphology, the colours of which are indistinguishable for E and S0 galaxies. The presence of such a large fraction of late-type galaxies on the CMR ridge line, with no increase in the CMR scatter, is surprising. However, Fig. 8 shows that our 13-arcsec aperture is dominated by bulge light, and that in every case the galaxies possess only a very faint disc. This could be a low-redshift analogue of the trend seen in high-redshift clusters by Dressler et al. (1997), who conclude that ellipticals predate the cluster virialization, but that late-type galaxies turn into S0 galaxies upon encountering the cluster. These anaemic spirals are likely to be galaxies in which star formation has been suppressed by ram-pressure stripping, but which still retain late-type characteristics (Poggianti et al. 1999).

Using the photometric diameter $D_V$ we have compared the colour magnitude relation found within a fixed 13-arcsec diameter and that measured within $D_V$. As expected, that measured within $D_V$ is shallower, owing to colour gradients within the galaxies. The slope of the relation is still easily seen, however, showing that the intrinsic colours of galaxies vary systematically as a function of magnitude.

The slope of the CMR also remains constant as a function of radius within cluster, as expected from a universal metallicity–mass relation. We subtract the mean relation from the colours in order to study the gradients within the cluster. We find evidence for a gradient in the CMR-corrected colours with projected cluster radius. Using a naive calculation for the expected slope in the photometric zero-points, we conclude that it is at least a 6σ result. The slope of the gradient is approximately one third of the size of that found by Abraham et al. (1996) in the $z = 0.23$ cluster Abell 2390, who attributed this to a gradient in the mean ages of the galaxies. We also find some evidence for decreasing scatter in the early-type galaxies towards the outskirts of the cluster as compared with the central parts. The upper limit for $E(U - V)$ (Ferguson 1993) in the cluster core is approximately equal to the $U - V$ colour gradient observed; however, we calculated the increased scatter produced in the CMR by the presence of enough dust in the cluster core to account for the entire gradient, and found that it was greater than the scatter observed in the elliptical galaxies. We therefore conclude that there cannot be sufficient dust in the cluster core to account for the entire gradient, and at least some of this gradient must be caused by a systematic variation in galaxy age.

By comparing the colours with the stellar evolution models of Bower et al. (1998), we estimate that the maximum age difference between the galaxies in the center of the cluster and in the outskirts is 2 Gyr. Although younger galaxies show more scatter in the CMR than old galaxies (BLE92b; Bower et al. 1998), the core of the cluster is contaminated by young galaxies in front of and behind the cluster.

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APPENDIX A: SEEING CORRECTIONS

In this discussion, we restrict ourselves to a circularly symmetric point spread function (PSF). All our measured properties are circularly averaged, so any non-spherical symmetry in the PSF, resulting maybe from poor tracking or focus, would cause only second-order effects (Saglia et al. 1993a).

The Fourier transform of the PSF can be predicted using atmospheric turbulence theory to be $\exp[-(kb)^{5/3}]$ (Fried 1966; Woolf 1982), where the scaling parameter $b = \text{FWHM}/2.9207006$. We generalize this to

$$p_g(k) = \exp[-(kb)^{\gamma}], \quad (A1)$$

where $\gamma$ controls the amount of light in the wings of the PSF. $\gamma = 2$ corresponds to a Gaussian profile, while the theoretically predicted value of $\gamma = 5/3$ gives a more wingy PSF. Lower values of $\gamma$ produce even larger wings (e.g. Saglia et al. 1993a).

We used least-squares fits of the brightest stellar objects in each image to obtain both $\gamma$ and $b$ for each exposure. Fig. A1 shows the distribution of FWHM for all of our images. Although the values of $\gamma$ vary from 1.2 to 1.8 (although mainly clustered around 1.45), there is very little variation between stars in the same exposure, so when calculating the final seeing correction for a galaxy we use the $\gamma$ corresponding to the image the galaxy was measured from.

The intensity at a radius $R$ from a source $I(R)$ is then given by the convolution of the surface brightness distribution of the object in the sky $[I^*(R)]$ with the PSF $[p_g(R)]$:

$$I(R) = I^*(R) \otimes p_g(R). \quad (A2)$$

For stellar objects, we simply take the intensity distribution on the sky to be a delta function, and for galaxies we use the canonical deVaucouleurs $R^{1/4}$ law:

$$I^*(R) = \left\{ \begin{array}{ll}
I_e \exp \left\{ -7.669 \left( \frac{R}{R_e} \right)^{1/4} - 1 \right\}, & \text{galaxy}, \\
L_0 \delta_d(R), & \text{star},
\end{array} \right. \quad (A3)$$

where $R_e$ is the half-light radius, at which $F(R_e) = F(\infty)/2$, $I_e = I(R_e)$, $\delta_d(R)$ is the Dirac delta and $L$ is the luminosity of the star.

To convert the luminosity distributions of an object in the sky into the luminosity distribution of the object on the detector we must convolve with the PSF (equation A2).

To seeing-correct our objects we require the difference, in

![Figure A1. The distributions of the seeing FWHM obtained by least-squares fitting of the profile of the brightest stars in each image to the theoretical seeing PSF described in equation (A1).](image)

**Table 8.** Table of galaxy seeing corrections based on a galaxy with $R_e = 5$ arcsec. To obtain the seeing-corrected value, the seeing correction is subtracted from the observed aperture magnitude. We parametrize the profile of the PSF using $\gamma = 1.47$, the average $\gamma$ for our observations. Corrections are shown for circular apertures ranging from 4 to 60 arcsec diameter. When seeing-correcting our data, we calculated a correction for each image by fitting a FWHM and $\gamma$ to it from the bright stars.

| seeing | 4   | 6   | 8   | 10  | 13  | 16  | 20  | 25  | 32  | 40  | 50  | 60  |
|--------|-----|-----|-----|-----|-----|-----|-----|-----|-----|-----|-----|-----|
| FWHM  |     |     |     |     |     |     |     |     |     |     |     |     |
| 4.50  | 0.909 | 0.564 | 0.376 | 0.268 | 0.176 | 0.126 | 0.087 | 0.060 | 0.039 | 0.027 | 0.018 | 0.014 |
| 4.00  | 0.789 | 0.476 | 0.314 | 0.222 | 0.146 | 0.104 | 0.072 | 0.050 | 0.033 | 0.022 | 0.015 | 0.011 |
| 3.50  | 0.667 | 0.391 | 0.254 | 0.180 | 0.118 | 0.085 | 0.059 | 0.041 | 0.027 | 0.018 | 0.012 | 0.009 |
| 3.00  | 0.542 | 0.309 | 0.200 | 0.141 | 0.093 | 0.067 | 0.046 | 0.032 | 0.021 | 0.014 | 0.010 | 0.007 |
| 2.50  | 0.418 | 0.233 | 0.150 | 0.106 | 0.070 | 0.050 | 0.035 | 0.024 | 0.016 | 0.011 | 0.007 | 0.005 |
| 2.00  | 0.300 | 0.165 | 0.106 | 0.075 | 0.050 | 0.036 | 0.025 | 0.017 | 0.011 | 0.008 | 0.005 | 0.004 |
| 1.50  | 0.193 | 0.106 | 0.068 | 0.049 | 0.032 | 0.023 | 0.016 | 0.011 | 0.007 | 0.005 | 0.003 | 0.002 |
| 1.00  | 0.104 | 0.057 | 0.037 | 0.027 | 0.018 | 0.013 | 0.009 | 0.006 | 0.004 | 0.002 | 0.001 | 0.001 |

**Table 9.** Table of stellar seeing corrections. To obtain the seeing-corrected value, the seeing correction is subtracted from the observed aperture magnitude. We parametrize the profile of the PSF using $\gamma = 1.47$, the average $\gamma$ for our observations. Corrections are shown for circular apertures ranging from 4 to 60 arcsec diameter. When seeing-correcting our data, we calculated a correction for each image by fitting a FWHM and $\gamma$ to it from the bright stars.

| seeing | 4   | 6   | 8   | 10  | 13  | 16  | 20  | 25  | 32  | 40  | 50  | 60  |
|--------|-----|-----|-----|-----|-----|-----|-----|-----|-----|-----|-----|-----|
| FWHM  |     |     |     |     |     |     |     |     |     |     |     |     |
| 4.50  | 1.093 | 0.573 | 0.332 | 0.214 | 0.129 | 0.088 | 0.060 | 0.041 | 0.028 | 0.020 | 0.014 | 0.011 |
| 4.00  | 0.922 | 0.461 | 0.263 | 0.170 | 0.104 | 0.072 | 0.049 | 0.034 | 0.023 | 0.016 | 0.012 | 0.009 |
| 3.50  | 0.746 | 0.357 | 0.202 | 0.132 | 0.082 | 0.057 | 0.040 | 0.028 | 0.019 | 0.013 | 0.010 | 0.007 |
| 3.00  | 0.573 | 0.263 | 0.150 | 0.099 | 0.063 | 0.044 | 0.031 | 0.022 | 0.015 | 0.011 | 0.008 | 0.006 |
| 2.50  | 0.408 | 0.184 | 0.107 | 0.072 | 0.046 | 0.033 | 0.023 | 0.016 | 0.011 | 0.008 | 0.006 | 0.004 |
| 2.00  | 0.263 | 0.120 | 0.072 | 0.049 | 0.032 | 0.023 | 0.016 | 0.012 | 0.008 | 0.006 | 0.004 | 0.003 |
| 1.50  | 0.150 | 0.072 | 0.044 | 0.031 | 0.020 | 0.015 | 0.011 | 0.008 | 0.005 | 0.003 | 0.002 | 0.002 |
| 1.00  | 0.072 | 0.037 | 0.023 | 0.016 | 0.011 | 0.008 | 0.006 | 0.004 | 0.003 | 0.002 | 0.001 | 0.001 |
magnitudes, of the flux of an object as measured within an aperture of radius \( R \) on our detector \([F(R)]\) and its flux as measured within the same aperture on the sky \([F' (R)]\). To find the flux inside an aperture of radius \( R \), we simply integrate the required luminosity distribution.

We found the value of \( R \) to give little effect on the seeing correction for all \( R \leq 1 \), so we use \( R = 5 \) arcsec for all of our galaxies.

Numerical integration techniques were used to perform both the integrations and the convolutions, the results of which for a variety of apertures and seeing FWHM are shown in Tables 8 and 9. Both of these tables assume \( \gamma = 1.47 \), the average value for our observations.

APPENDIX B. PHOTOMETRIC CALIBRATION VIA FRAME OFFSETS

The method we used to generate this system is very similar to that used in Maddox et al. (1990) to homogenise the APM galaxy catalogue. However, it is far simpler, owing to our detector’s better flat fielding and its linearity.

To calculate a set of zero-point offsets for each image the regions of overlap between each pair of images was examined. An object positioned in an area where frames \( i \) and \( j \) overlap will have magnitudes \( m_i \) and \( m_j \) as measured from frames \( i \) and \( j \) respectively. The actual magnitude for this object is \( m_0 \), and in the absence of observational errors these three quantities can be related to each other thus:

\[
m_0 = m_i + C_i = m_j + C_j, \tag{B1}
\]

where \( C_i \) is the correction applied to the zero-point of image \( i \). If we define

\[
T_{ij} = m_i - m_j = C_j - C_i
\]

as the overlap difference between \( i \) and \( j \), then the offset correction for any image can be calculated from the overlap differences and offset corrections of any adjacent overlapping image.

\[
C_j = C_i + T_{ij}, \forall j \subset i,
\]

where \( j \subset i \) denotes any pair of images \( i,j \) with a valid overlap region, and \( C_j \) denotes the \( C \) as calculated from image \( j \). First we construct an observed estimate of the \( T_{ij} \),

\[
T_{ij}^n = \frac{\sum (m_i^n - m_j^n)}{N_{ij}},
\]

where \( N_{ij} \) is the number of bright objects in the region of overlap between images \( i \) and \( j \), and the \( m_i^n \) and \( m_j^n \) are the measured magnitudes of object \( n \) in images \( i \) and \( j \) respectively.

Now we can use a weighted mean of the \( T_{ij} \) to find a value for the photometric offsets.

\[
C_{i}^{n+1} = \left( \frac{C_i^n W_i + \sum_{j \subset i} (C_j^n + T_{ij}^n) W_{ij}}{W_i + \sum_{j \subset i} W_{ij}} \right), \tag{B2}
\]

where the \( C_i^n \) are the \( C \) calculated in iteration \( n \), and \( W_j \) is the mean of the weights, \( W_{ij} \). These weights were chosen to be proportional to the number of objects used to calculate the \( T_{ij}^n \) and normalised to be in the range 0 to 1. We chose to use the number in the overlap \( i,j \), rather than the inverse of the scatter in the \( (m_i^n - m_j^n) \). With the low number of objects present in our overlaps, the scatter is not always well determined, and can be artificially low for overlaps with low \( N_{ij} \), just the opposite of the required behaviour.

This iterative method does not constrain the total photometric offset, it merely ensures the best possible relative photometry of the system by removing as much of the drift in the zero-point between images as is possible in a self consistent manner. We therefore arbitrarily renormalize the \( C_i \) after every iteration so that they have zero mean.

Equation (B3) was iterated to find the best set of \( C_i \). To measure the progress of the iterations, we construct a measure of the homogeneity of the system after iteration \( n \),

\[
E^n = \sum_{i,j} W_{ij} (T_{ij} + C_i^n - C_j^n)^2, \tag{B3}
\]

where \( N \) is the total number of images for which we are trying to ascertain the \( C_8 \). We iterate until the rate of change of \( E^n \) has slowed to less than \( E^n/1000 \) per iteration.

Obviously much care has to be taken in the measurement of the \( T_{ij} \). We must ensure that effects such as different seeing conditions on the two overlapping images do not cause any systematic offsets. Although seeing conditions are taken into account for each object when measuring the \( m_i \) (see Section 6), the effects of seeing on objects adjacent to the aperture are not corrected for, i.e. more light from an adjacent object will enter the aperture for images with worse seeing. We therefore measured the \( m_i \) in various sized apertures, from 8.8 to 26 arcsec diameter, and using background annuli from 20 to 100 arcsec diameter with 25 and 3.2 arcsec widths. If the \( T_{ij} \) are well behaved under all measuring conditions, then we simply take the median value. If they are not well behaved, we inspect the region more carefully to ascertain the cause, or we mark the \( T_{ij} \) as unreliable.

To summarize, we can now construct a homogeneous data set by adding to each magnitude \( m_i \) the offset correction for the image in which it was measured \( (C) \) to obtain the object’s corrected magnitude \( m_0 \) (see equation 4). This homogeneous photometric system, however, is only relative. To set the overall photometric offset, we use the fixed aperture magnitudes of BLE92a. We have new observations of all of the BLE92a sample, so using these galaxies we can define a transformation between our corrected aperture magnitudes and theirs. We conclude that our V-band response is similar to that of BLE92a, but the larger value of the U-band colour term indicates that our U-band response is not such a good match, probably owing to our increased sensitivity to the blue portion of the U bandpass.

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