The Halo Stars in NGC 5128. III:
An Inner-Halo Field and the Metallicity Distribution

William E. Harris

Department of Physics & Astronomy, McMaster University, Hamilton ON L8S 4M1

harris@physics.mcmaster.ca

and

Gretchen L. H. Harris

Department of Physics, University of Waterloo, Waterloo ON N2L 3G1

glharris@astro.uwaterloo.ca

ABSTRACT

We present new Hubble Space Telescope WFPC2 (V, I) photometry for field stars in NGC 5128 at a projected distance of 8 kpc from the galaxy center, which probe a mixture of its inner halo and outer bulge. The color-magnitude diagram shows an old red-giant branch which is even broader in color than our two previously studied outer-halo fields (at 21 and 31 kpc), with significant numbers of stars extending to Solar metallicity and higher. The peak frequency of the metallicity distribution function (MDF) is at \([m/H] \simeq -0.4\), with even fewer metal-poor stars than in the outer-halo fields. If we use the 21- and 31-kpc fields to define template “halo” MDFs and subtract these from the 8-kpc field, the residual “bulge” population has a mean \([m/H] \simeq -0.2\), similar to the bulges of other large spirals and ellipticals. We find that the main features of the halo MDF can be reproduced by a simple chemical evolution model in which early star formation goes on simultaneously with an initial stage of rapid infall of very metal-poor gas, after which the infall dies away exponentially. Finally, by comparison with the MDFs for the NGC 5128 globular clusters, we find that in all the halo fields we have studied, there is a clear decrease of specific frequency \(S_N\) (number of clusters per unit halo light) with increasing metallicity. At the lowest-metallicity range ([Fe/H] \(< -1.6\)) \(S_N\) is \(\sim 4 - 8\), while at metallicities [Fe/H] \(> -1\) it has dropped to \(\simeq 1.5\).
This trend may indicate that globular cluster formation efficiency is a strong function of the metallicity of the protocluster gas. However, we suggest an alternate possibility, which is that globular clusters form preferentially sooner than field stars. If most of the cluster formation within a host giant molecular cloud (GMC) takes place sooner than most of the distributed field-star formation, and if the earliest, most metal-poor star-forming clouds are prematurely disrupted by their own first bursts of star formation, then they would leave relatively few field stars with a high–$S_N$ population. The high specific frequency at low metallicity may therefore be related to the comparably large $S_N$ values found in the most metal-poor dwarf ellipticals.

Subject headings: galaxies: halos — galaxies: individual (NGC 5128) — galaxies: stellar content — galaxies: formation

1. Introduction

An important part of the historical record of galaxy formation is locked up in the ages and metallicities of the earliest stellar populations, particularly those found in the halos of galaxies. For most galaxies, however, even the brightest old-halo stars are too dim to be seen as anything but a smooth distribution of integrated light, from which it is difficult to deduce much more than the mean age and metallicity of the population (e.g., Trager et al. 2000). Giant ellipticals present a special challenge because they are the only galaxy type not found in the Local Group, thus few are close enough to allow direct, comprehensive probes into their stellar populations.

NGC 5128, the central giant elliptical in the Centaurus group just 4 Mpc distant, is an outstanding exception. Recently, we used deep ($V, I$) photometry from $HST/WFPC2$ to obtain color-magnitude diagrams for fields in the outer halo of NGC 5128, at projected galactocentric distances of 21 and 31 kpc (G.Harris et al. 1999; G.Harris & W.Harris 2000, hereafter Papers I and II). The CMDs of these halo fields reveal the top $\sim 2.5$ magnitudes of a normal old red-giant branch (RGB), with a wide range in color that is indicative of a broad metallicity range. Since the RGB loci for old stars are predominantly sensitive to metallicity rather than age, they can be used to construct a first-order metallicity distribution function (MDF) – the first such one for any giant elliptical and one which does not rely on indirect arguments from luminosity-weighted integrated light. The MDF is an important step towards constructing an eventual age-metallicity relation (AMR) to constrain its evolutionary history.

In the present study, we present similar data for an “inner halo” (or alternatively an “outer bulge”) field much closer in to the center of NGC 5128, at a projected distance of 8 kpc. This field reveals important differences in the MDF compared with the outer-halo fields and, perhaps, give us the first direct look at the pure bulge population of a giant elliptical.
2. Observations and Data Analysis

The location of our 8-kpc target field is southwest of the center of NGC 5128 with the PC1 chip center at 13\(^h\)24\(^m\)53\(^s\), −43\(^o\)04′35″ (J2000). This location avoids the central dust lane (see Figure 1 of Paper II for a finder chart), and comparison with a deep \((B − R)\) color image by Peng (2001) sensitive to the presence of dust shows that it also avoids two thin dust streamers extending southward along the major axis. The raw observations, taken in HST Cycle 8 for program GO-8195, consisted of a sequence of 14 half-orbit exposures in F606W (wide \(V\)) totalling 17500 seconds, and 10 exposures in F814W (wide \(I\)) totalling 12100 seconds, the same as for our 31-kpc field taken during the same program.

Data analysis paralleled what we did for the 31-kpc field and our procedures are discussed in detail in Paper II. Pointings in the sequence of raw images differed by several pixels in \((\Delta x, \Delta y)\), and we reregistered the individual exposures and median-combined them with the normal IRAF and STSDAS packages to produce clean median images free of cosmic rays, bad pixels, and other artifacts. Photometry on the master \(V\) and \(I\) frames was then carried out with the DAOPHOT II code (Stetson 1992) in the same manner as described in Papers I and II, with a two-pass sequence of FIND, PHOT, and ALLSTAR. Transformation of the instrumental magnitudes to final \(V, I\) values followed the prescriptions of Holtzman et al. (1995).

The one important respect in which the data reduction for this field differed from our previous work arose from the level of stellar crowding. While the two outer fields (Papers I, II) are completely uncrowded at the characteristic WFPC2 0″1 resolution, the 8-kpc field has a sufficiently high density of resolved stars to make the photometry in the three WF chips quite difficult and any discussion of these is not pursued further here. On the PC1 chip with its superior resolution, the photometry is much more easily manageable with careful attention to defining the stellar point spread function (PSF). We carried out complete reductions of the PC1 images with five different empirically constructed PSFs: (1) the Gaussian-plus-residuals model in DAOPHOT, constructed from an average of several bright stars in the PC1 chip with three iterations of “cleaning” neighbors from around the PSF candidate stars, (2) the Lorentzian-plus-residuals model in DAOPHOT, (3) PSFs for \(V\) and \(I\) supplied from the STScI archive, (4) PSFs derived from uncrowded bright stars on other PC1 images (Stetson, private communication), and (5) PSFs taken from the PC1 images in our uncrowded outer-halo fields. In each case we independently rederived the necessary aperture corrections through curves of growth to normalize the PSF-fitted magnitudes to the 0″5 aperture radius used in the Holtzman et al. calibrations.

The results which clearly gave the cleanest fits to the image, and which we adopted, were the Lorentzian-based PSFs constructed iteratively from the 8-kpc PC1 field itself. However, despite differences in the PSF profile shapes and the consequent aperture corrections, the final lists of calibrated \((V, I)\) stellar magnitudes (essentially resting on the PSF scalings derived by ALLSTAR) proved to be very similar across all five types of PSF models.

The final color-magnitude diagram (CMD) in \((I, V − I)\) for a total of 17,283 stars on PC1
detected in both filters is shown in Figure 1. The full dataset can be obtained from the authors on request. For this inner field, we made no attempt to find or remove nonstellar objects from the list; although there are a few faint background galaxies visible in the field, their numbers (from the image classification analysis done more completely in Papers I and II) are quite small here compared with the overwhelmingly large population of stars, and the high degree of crowding would, in any case, vitiate objective measurements of nonstellar image shapes for most such objects.

The CMD, when compared with our previous data for the outer-halo fields, shows the same old RGB population with a bright-end “tip” at \(I \simeq 24.2\), but extends even further to the red than did the outer fields. The number of stars appearing above the RGB tip \((I < 24)\) that might arise from a younger AGB population, for example (see Soria et al. 1996, and Papers I and II), is statistically negligible. Within the main RGB distribution, however, our deductions about the range of metallicities of these stars depend on the effects of photometric measurement scatter, and the much higher level of crowding here makes this question more acute. To investigate the photometric scatter, we carried out several simulations in which 500 stars at a time were added to the original images, with input colors and magnitudes following an ideal, narrow RGB sequence. These images were then measured in exactly the same procedure (two-pass FIND/ALLSTAR), yielding the results shown in Figure 2.

The artificial-star tests show higher scatter than did the same tests on the outer fields (as expected from the higher degree of crowding; compare Figure 5 from Paper II), but within the upper part of the RGB, the observed color spread is far larger than can be explained by photometric errors and must be intrinsic to the stellar population. The trend of measurement scatter in each band is shown in Figure 3. Encouragingly, in both \(V\) and \(I\) the simulations revealed no photometric bias larger than \(~0.03\) mag at any level of interest here, i.e. much smaller than the level of measurement uncertainty. We therefore made no attempts to apply any bias corrections.

Lastly, the artificial-star tests were used to estimate the photometric detection completeness as a function of magnitude. The levels at which the numbers of detected stars drop to 50% of the numbers put in can be expressed (once the instrumental \((v, i)\) magnitudes are folded through the Holtzman et al. calibration relations) as

\[
I_{50} = 26.84 - 0.062(V - I) + 0.025(V - I)^2 \quad \text{for } I,
\]

\[
I_{50} = 28.76 - 1.247(V - I) + 0.065(V - I)^2 \quad \text{for } V.
\]

Approximate completeness curves in \(V\) and \(I\) are shown in Figure 4. The two smooth curves are Pritchet interpolation functions (Fleming et al. 1995) of the form

\[
f = 0.5 \left( 1 + \frac{a(m - m_0)}{\sqrt{1 + a^2(m - m_0)^2}} \right)
\]

where \(m_0\) is the magnitude at which 50% completeness is reached and \(a\) is a parameter governing the steepness of the decline in \(f\). The curves shown in the Figure are for a particular color \((V - I) = 2\).
and are only to be taken as indicative of the average magnitudes at which the photometry becomes severely incomplete. In practice, we define $f$ in terms of the instrumental magnitudes $(v, i)$ and then translate them through the photometric calibration equations into the appropriate $f(V, I)$ for each point in the CMD.

For the purposes of deriving the stellar metallicity distribution, detection incompleteness is unimportant except at the extreme upper right corner of the CMD, where the very reddest and brightest RGB stars fall near the completeness cutoff determined by the $V$ filter. As will be seen below, these reddest stars are likely to have above-Solar heavy-element abundances; they are redder than the old-RGB populations in any previously observed globular clusters, dwarf galaxies, or halos of larger galaxies. Despite our deliberately planned exposure times in $V$, which were considerably longer than in $I$ (and our usage of the “wide $V$” F606W filter rather than the more nearly standard F555W), we may still be missing some fraction of the very reddest RGB stars. In the discussion below on the metallicity distribution, we explicitly include the effects of incompleteness as far as the data allow. However, it should be kept in mind that if the outer bulge of NGC 5128 does contain stars well above Solar metallicity, our data are not capable of detecting most of them. Such objects, if they exist, will have to be searched for with longer-wavelength $JHK$ imaging, such as with the NICMOS camera (see Marleau et al. 2000).

3. The Metallicity Distribution Function

3.1. Calibrating the Metallicity Measurement

The derivation of stellar metallicities follows the procedure developed in Paper II: briefly, we superimpose a stellar model grid on the CMD and interpolate within it to calculate a heavy-element abundance $Z$ for each star. Our adopted metallicity index is then by definition $[m/H] = \log \left( \frac{Z}{Z_\odot} \right)$. The main stellar model grid is that of VandenBerg et al. (2000), which provides fiducial RGB evolutionary tracks in metallicity intervals of roughly 0.1 dex in $\log Z$ from $Z = 0.00017$ up to $Z = 0.013$. Solar metallicity is adopted as $Z_\odot = 0.017$. However, a significant fraction of the stars in our inner-halo sample are more metal-rich than this, and particular attention needs to be paid to extending the grid to higher $Z$. Unfortunately, it is precisely this high-$Z$ regime which is fraught with the greatest uncertainty in every direction. In Paper II, we used two metal-rich isochrone lines from Bertelli et al. (1994) ($Z = 0.02, 0.05$) to extend the grid. However, these have scaled-Solar element ratios and are thus fundamentally not compatible with contemporary $\alpha-$enhanced models. This issue had little relevance for the outer-halo fields (Papers I and II) because these fields did not have many stars which extended into the high-$Z$ range, but for our inner-halo data the question is more pressing. Stars with $[\alpha/Fe] > 0$ are hotter than ones with scaled-Solar ratios because of their lower opacities, and simple renormalization of the total $Z$ becomes risky particularly at high metallicity (VandenBerg et al. 2000; Salaris & Weiss 1998).

Ideally, we would like to employ a grid of RGB models which is (a) computed with full
\(\alpha\)-enhancement, (b) extends to at least \(3Z_\odot \simeq 0.05\), (c) is densely spaced in \(\log Z\), and (d) also properly accounts for any progressive change in \([\alpha/Fe]\) with metallicity itself (e.g. Shetrone et al. 2001). Such material is not yet available. Recently Salasnich et al. (2000) have published metal-rich and \(\alpha\)-enhanced isochrones for a widely spaced grid of heavy-element values \(Z = (0.008, 0.019, 0.04, 0.07)\). However, at least in the upper-RGB luminosity range which we require, the theoretical \((V - I)\) colors of all these tracks are significantly bluer for a given metallicity compared with the VandenBerg et al. (2000) models, as well as (more importantly) the real star clusters that we use to calibrate the grid (see below). A more finely spaced grid of metal-rich models is available from Yi et al. (2001), but these are calculated for scaled-Solar abundances. For the present we are forced to adopt an alternate procedure.

Our approach is a continuation of the one we adopted in Paper II: instead of relying solely on theoretically computed color indices, we pin the model tracks as closely as possible to the RGB loci and abundances of real Milky Way clusters. At low \(Z\), this method works handily because there are many well observed globular clusters with low metallicities, low reddenings, and accurate RGB loci in the \((M_I, V - I)\) plane. The five which we use to cover the range from \([m/H] = -2.0\) up to \([m/H] = -0.4\) are M15, NGC 6397, NGC 6752, NGC 1851, and NGC 104 (47 Tuc). At higher \(Z\), the situation is considerably worse. The two metal-rich objects we rely on, out of sheer necessity, are the bulge globular cluster NGC 6553 and the very old disk cluster NGC 6791. NGC 6553 is generally thought to have essentially Solar heavy-element abundance (Barbuy et al. 1999; Cohen et al. 1999; Carretta et al. 2001; Beaulieu et al. 2001; Origlia et al. 2001) with \([Fe/H] \simeq -0.2 \pm 0.1\) and \([\alpha/Fe] \simeq 0.2 - 0.3\). NGC 6791 is generally thought to have \([Fe/H] \simeq +0.4\) with little or no \(\alpha\)-enhancement (Peterson & Green 1998; Chaboyer et al. 1999), but see Taylor (2001) for a healthy discussion of concerns. Both clusters have observational handicaps that are much more serious than in the low–\(Z\) regime: NGC 6553 is both heavily reddened and differentially reddened, while NGC 6791 has an age which is younger than the standard globular clusters by \(\sim 2\) Gy (e.g. Chaboyer et al. 1999). To work around these difficulties, we have taken the Yale scaled-Solar tracks (Yi et al. 2001) and have used them to define the shapes of the tracks in the \((M_I, V - I)\) plane near \(Z = Z_\odot\).3 For NGC 6791, we adopt a fiducial value \(Z = 0.045\) and use the models to find that we must shift its upper RGB 0.15 mag redder in \((V - I)_0\) to correct it to an age 2 Gy older. This adjusted locus, by definition, is the uppermost boundary of the metallicity sequence we can use at present. Lastly, we also interpolate within the models to define a track at \(Z = 0.025\), in between NGC 6553 and NGC 6791.

The complete array of fiducial tracks and star clusters that we have adopted is shown in Figure 5. The fiducial points for NGC 6553, for which \(Z \simeq 0.017 = Z_\odot\), fall slightly redder than the most metal-rich VandenBerg et al. track at \(Z = 0.013\), and the raw NGC 6791 sequence (not corrected for age differences) lies as it should between the tracks for \(Z = 0.025\) and 0.045. It is obvious

---

3The VandenBerg et al. (2000) models use \(Z_\odot = 0.017\) while the Yale models use \(Z_\odot = 0.018\). At the level of accuracy needed for our purposes, this difference is negligible.
that these observational constraints on the high-$Z$ part of the grid, from NGC 104 up to NGC 6791, are very thin indeed and that our current definition of this part of the grid is only a stopgap measure. With further careful observations of selected globular clusters, it may be possible to add one or two more fiducial sequences in this range, but particularly urgent attention is needed to the computation of appropriate models.

We emphasize once again that the extension of this interpolation grid above Solar metallicity is the most uncertain part of our data analysis.

3.2. Results for the MDF

Once the model grid is defined, we interpolate linearly within it, in the manner described in Paper II, to estimate the metallicity of each star. Compared with the previous analysis of Paper II, we have modified the interpolation code to add minor improvements in the calculation of bolometric correction and the ability of the code to use stars closer to the RGB tip. For purposes of homogeneity, we have therefore recalculated the MDFs for all three of our program fields (8, 21, and 31 kpc) through the same updated routine.

As part of the interpolation routine we also compute an uncertainty $\epsilon(Z)$ in the abundance for each star, generated by the random photometric error in magnitude and color at its position in the CMD. These uncertainties are determined numerically by simply adding $\epsilon(I)$ and $\epsilon(V - I)$ to the given $(I, V - I)$ for the star and recalculating $Z$. In practice, the size of $\epsilon(Z)$ is determined almost entirely by the scatter in color, $\epsilon(V - I)$, because the RGB evolutionary tracks are nearly vertical in the bolometric CMD and so small shifts in magnitude alone do not affect the resulting MDF.

The data for our 8-kpc field, superimposed on the RGB model tracks, are redisplayed in Figure 6 in the $(M_{bol}, (V - I)_0)$ plane within which the interpolation process is actually carried out. The resulting histograms in $[\text{m/H}]$, divided somewhat arbitrarily into three luminosity bins, are shown in Figures 7 and 8. Revisiting the combined MDF for the two outer fields (Figure 7), we see in all bins a broad peak near $[\text{m/H}] \simeq -0.5$, a steep decrease toward higher metallicity, and a longer tail toward lower metallicity. There are nearly negligible numbers of stars either more metal-poor than $[\text{m/H}] \lesssim -1.5$ (which is, notably, the metallicity range including most of the halo stars in the Milky Way) or more metal-rich than $[\text{m/H}] \gtrsim -0.2$. In the fainter luminosity bins, the MDF becomes more spread-out due to the increasing photometric scatter, but the peak location and general distribution shape remain the same at all magnitudes.

For our 8-kpc field (Figure 8) the patterns are not quite the same. As we expected from the raw CMD, the populations in all the bins have higher proportions of metal-rich stars compared with their outer-halo counterparts, but the fainter bins shift progressively further to the metal-rich side. This trend apparently violates the requirement of internal self-consistency, i.e. that the MDF for the same population of stars should be independent of the luminosity range sampled. However, the highly consistent results from the outer fields (Figure 7) indicate that the problem is not an
artifact of the interpolation routine or the model grid itself. A large part of the explanation is likely
to be the detection incompleteness for extremely red stars, which will affect the highest-luminosity
bins somewhat more because the RGB stars there (see Fig. 1) lie closer to the completeness cutoff
line than do equally metal-rich stars a magnitude further down the giant branch.

To take partial account of incompleteness effects, we recalculate completeness-corrected MDFs
for both the inner and outer fields (where now each star is counted as \((f_V f_I)^{-1}\), the inverse of
the photometric completeness at that point on the CMD; to avoid wild excursions in the MDF
we arbitrarily set the maximum correction factor at 2.0). From Figure 8, it is evident that the
metal-richer side of the distribution is enhanced as a result, and more so in the higher-luminosity
bins. For the outer halo, incompleteness is much less important because the RGB does not reach
far enough to the red to be strongly affected. For \(f\)-values lower than \(\simeq 0.5\), completeness effects
become poorly understood and very risky to apply. We regard the corrected histograms in Figure 8
only as an informed guess at the red-end shape of the MDF, and not a substitute for raw data that
reach appropriately deeper in the first place.

Lastly, we note that the fainter bins become progressively more affected not only by photo-
metric spread but also by unrecoverable systematic errors: higher and higher fractions of the bluer
stars are driven by photometric scatter out of the model grid entirely, but the stars driven to the
red side by the same photometric random scatter stay within the grid and skew the resulting MDF
to higher metallicity. See Paper II for more discussion. In what follows, we restrict our discussion
to the brightest interval \(M_{bol} < -2.5\) which is least affected by this combination of errors and
uncertainties.

Our final completeness-corrected histograms for the inner and outer fields are summarized in
Table 1 and shown in Figure 9. Regardless of the level of completeness corrections, the inner-field
MDF is clearly more metal-rich on average, and extends to higher maximum metallicity, than the
outer halo.

### 3.3. The MDF as a Probability Distribution

As we discussed in Paper II, a physically natural way to display the MDF which is easily con-
ected to chemical evolution models, is as the number of stars per unit heavy-element abundance,
\(dn/dZ\). The numbers in Table 1 can readily be converted to this alternate form, but while we are
at it, we can also partially compensate for the effects of random photometric errors. Define \(p(Z)\)
as the probability of finding stars at a given abundance \(Z\) as follows:

\[
p(Z) = P_0 \sum_{i=1}^{n} \frac{1}{f_i} \cdot e^{-(Z-Z_i)^2/2\epsilon_i^2}.
\]  

(1)

In this expression, each star is replaced by a small Gaussian centered at its measured abundance
\(Z_i\) and with dispersion \(\epsilon_i\) equal to the uncertainty in its calculated heavy-element abundance \(Z\).
As before, $f$ is the detection incompleteness at the star’s location in the CMD. The normalization constant $P_0$ is chosen to set $\int p(Z)dZ \equiv 1$ over all $Z$.

The probability distributions $p(Z)dZ$ for the inner and outer fields are shown in Figure 10. The curve for the outer fields is smooth and featureless, reaching a peak at $Z(\text{max}) \approx 0.25Z_{\odot}$ (compare Fig. 13 from Paper II, where basically the same material is plotted in raw histogram form). By contrast, the curve for the inner field is extremely broad, declining slowly from a peak at $\approx 0.35Z_{\odot}$ toward levels that extend well above Solar abundance. We emphasize again, however, that its extension at high $Z$ is uncertain, as indicated by the difference between the completeness-corrected and uncorrected curves shown in Figure 10. We will use these distributions as the basis for some brief modelling discussions in section 4 below.

3.4. A First Look into the “Bulge” Population?

In any representation, the MDF for our 8 kpc field is plainly more extended, more metal-rich, and more complex than the outer-halo MDF. The inner MDF is likely to be broader because it is made up of a wider mixture of subcomponents from different parts of the galaxy. The effective radius of the bulge light is $r_e = 330''$ (van den Bergh 1976), corresponding to $\approx 6.4$ kpc. Our 8-kpc field is therefore only $\approx 1.3r_e$ from the center, and so a significant contribution from the inner bulge of the galaxy should be present in our MDF.

If we now assume that our inner-field MDF is made up of two major components (bulge and halo), and that the halo component is similar to the outer-field MDF, we can subtract one from the other to gain a first, albeit very rough, look at the MDF for the “pure bulge” component of NGC 5128. To normalize the outer-field population to the inner one, we use the metal-poor parts of the two MDFs where we can plausibly expect the halo components to dominate. In Figure 11, we plot the ratio of numbers of stars in each field more metal-poor than a given cutoff metallicity $[m/H]$. For any cutoff in the range $[m/H] \lesssim -1$, this ratio stays roughly constant at $n_{\text{in}}/n_{\text{out}} \approx 0.46 \pm 0.01$, indicating that the two MDFs match up well over that range. At higher metallicities, the proportion of inner-field stars rises steadily and the two MDFs no longer match one another.

In Figure 12, we show the result of subtracting one component from the other. The lower panel gives the residual, equal to $(n_{\text{in}} - 0.46 \cdot n_{\text{out}})$ (where both MDFs here are the completeness-corrected versions). The halo component ($[m/H] < -1$) has now subtracted cleanly away, and the remaining metal-rich part, by hypothesis, represents the stars belonging to the outer bulge of the galaxy. This residual MDF, rising steadily from $[m/H] \approx -0.8$ up to near-Solar abundance, shows a broad peak and roughly symmetric shape, but as indicated above, it probably still underestimates the numbers of stars with above-Solar metallicity because of the inadequately understood incompleteness effects of the photometry.

The formal mean and standard deviation of the residual “bulge” MDF as shown here are $\langle m/H \rangle = -0.19 \pm 0.01$ and $\sigma(m/H) = 0.25$ dex. For comparison, Trager et al. (2000) find that the
bulge metallicities for a large sample of other elliptical galaxies, determined from integrated-light spectral indices, lie in the typical range \([m/H] \simeq -0.1\) to \(+0.3\) with an average at \(+0.08\) (we quote here their determinations for the integrated light within \((r_e/2)\); their “core” metallicities within \((r_e/8)\) are slightly more metal-enhanced at \(\langle m/H \rangle = +0.26\). Given the presence of a modest metallicity gradient, our result for NGC 5128, measured further out in the bulge at \(1.3r_e\), appears to be quite consistent with their quoted range.

A comparison with the Milky Way bulge is also possible. The mean metallicity of the old Milky Way bulge stars within \(r_{gc} \sim 1\) kpc (corresponding to about half the classic effective radius) is at a very similar level: for example, Frogel et al. (1999) find \(\langle Fe/H \rangle = -0.19 \pm 0.03\) for five fields within \(3^\circ\) latitude (0.4 kpc) along the minor axis. McWilliam & Rich (1994) find \(\langle Fe/H \rangle = -0.27 \pm 0.12\) for 11 bulge giants at moderately low latitude, and Ibata & Gilmore (1995) find that the MDF for four Milky Way bulge fields peaks at \([Fe/H] \simeq -0.3\). Thus our tentative findings for the NGC 5128 “bulge” fall very much in line with the ranges observed in other elliptical and spiral bulges.

4. Modelling the MDF

The eventual aim of obtaining the metallicity distribution of the stellar population is to help place constraints on the evolutionary history of NGC 5128. Both theoretical modelling and recent observations suggest that giant ellipticals are likely to form or grow by a mixture of different processes including hierarchical merging of gas clouds in the early universe (e.g. Pearce et al. 1999; Cole et al. 2000; Dubinski 1998; Kauffmann et al. 1993; Theis et al. 1992; Forbes et al. 1997); merging of previously formed large disk galaxies (e.g. Barnes & Hernquist 1992; Hibbard & Yun 1999; Naab & Burkert 2001; Schweizer 1987; Ashman & Zepf 1992; Zepf et al. 2000); and nondissipative, hierarchical merging of stellar fragments (e.g. Côté et al. 1998, 2000, 2001). Which of these processes dominates for any particular gE almost certainly depends strongly on its environment and individual history.

Matching our data with some of these advanced modelling codes is far beyond the scope of the present paper. We will discuss a specific comparison with one of the semianalytical codes for hierarchical merging, GALFORM (Cole et al. 2000), in a later paper (Beasley et al. 2002). For the present, in this section we briefly discuss a much simpler chemical evolution model capable of generating the first-order features of the MDFs that we observe. The basic picture we adopt is one of hierarchical formation, in which the bulk of the stars in the halo are assumed to form at relatively early epochs within many gas-rich clouds that merged to build up the (eventual) giant elliptical. That is, the MDF is the result of a large number of individual star-forming events within these smaller clouds and within the growing potential well of the gE. This is almost the opposite view from what we adopted in the previous section, i.e. that the MDFs that we see in these three fields resulted from a single series of in situ star-forming events, rather than from two rather distinct subcomponents.
4.1. An Accreting-Box Model for the MDF

As we argued in Paper II, perhaps the most interesting feature of the MDF is the striking lack of low-metallicity stars in either the inner or outer halos. A classic closed-box model of chemical evolution with initial gas abundance $Z_0$ and nucleosynthetic effective yield $y$ leads to an MDF with a characteristic exponential-decay shape $dn/dZ \sim e^{-(Z-Z_0)/y}$. Without highly arbitrary adjustments to the assumed star formation rates or the IMF, these basic models do not produce the steep rise in the number of stars at low $Z$ (see again Fig. 10, and see Paper II and references cited there for more discussion of the alternatives).

An alternate and almost equally simple class of models is the “accreting-box” type, within which new gas is added to the “box” after each round of star formation (see, e.g., Larson 1972; Pagel & Patchett 1975; Binney & Merrifield 1998; Portinari et al. 1998). A more intuitive way to state this is simply to visualize that the many protogalactic clouds, which start with primordial unenriched gas, are all busily forming stars within their own small potential wells while they are simultaneously merging together. That is, the merger epoch does not wait until after the protogalactic clouds have completed their own local star formation, nor do the small clouds wait until they have merged before beginning star formation. To some degree, both steps must be happening at once.

To quantify this approach in a straightforward way, assume (following the notation of Binney & Merrifield 1998; Pagel & Patchett 1975) that the box has gas mass $M_g(t)$ and stellar mass $M_s(t)$ at any time $t$, and that star formation proceeds in discrete timesteps $\delta t$, within each of which a small fraction $e$ of the gas forms into stars. Of the mass $\delta M_s$ of newly formed stars, a fraction $\alpha$ stays locked up in stellar remnants and the remainder $(1-\alpha)$ is returned to the interstellar medium, partially enriched by the products of nucleosynthesis. As usual, the yield $y$ is defined as the amount of heavy elements released back into the ISM per unit remnant mass. By counting gains and losses in the mass $M_Z$ in heavy elements, we eventually find (Pagel & Patchett 1975; Binney & Merrifield 1998) that the net change of the gas mass in heavy elements after each step can be written as

$$ \delta M_Z = -Z \cdot \delta M_s + (1 - Z)y \cdot \delta M_s \simeq (y - Z) \cdot \delta M_s $$

where $\delta M_s = \alpha e M_g$ is the net gain in star mass. In general, there are no interesting analytic solutions to this relation (except the artificial case $M_g = \text{const}$), but it can readily be integrated numerically as soon as a prescription for the gas infall rate is given.

To specify a form for the gas accretion rate that will be both realistic and flexible enough to handle a variety of cases, we note that recent simulations of giant galaxy formation (e.g. Theis et al. 1992; Dubinski 1998; Pearce et al. 1999; Cole et al. 2000) indicate that large numbers of gaseous fragments can be expected to come inward onto the central galaxy over time periods of perhaps $\sim 2$ Gyr give or take factors of two, after which the infall rate gradually dies away. For our purposes, we parametrize this sequence as shown in Figure 13: the rate of mass infall is assumed constant for some initial period $\tau_1$, after which it falls off exponentially with an e-folding time $\tau_2$:

$$ \frac{\delta M}{\delta t} = k \quad (t < \tau_1), $$
An exponentially declining infall rate has also been used in the context of disk formation by (e.g.) Lacey & Fall (1983); Chiappini et al. (1997). This “delayed exponential” model has four defining parameters in addition to the basic yield rate $y$: these are the initial infall rate $k$ and timescale $\tau_1$; the decay time $\tau_2$; and the metallicity $Z_g$ of the infalling gas. After time $t$, the total mass in stars will be larger than the initial mass $M_0$ in the box by some factor $F \equiv M(t)/M_0$ which is determined by $k, \tau_1, \tau_2$. The final mass $M_f$ of the model can then be set equal to the observed mass of the present-day galaxy to estimate $M_0$, i.e. the initial “seed” mass of the proto-elliptical.

Although the step upward from the simpler closed-box model to an accreting-box model inevitably involves a large increase in parameter space, it is helpful that the different free parameters govern different parts of the predicted MDF, as shown schematically in Figure 14. The shape of the rising part of the MDF at low metallicity helps to fix $k$ and $\tau_1$; the intermediate-metallicity section is determined by $\tau_2$; and at late times and larger $Z$, the gas accretion has died away to negligible levels and the system approaches closed-box evolution fixed by the yield $y$. The true closed-box model is a special case of our accreting-box model where $k = 0$.

We assume, more or less arbitrarily, that the initial abundance of the gas is primordial, $Z(0) = 0$. We then adopt an abundance $Z_g$ for the infalling gas and vary $k, \tau_1,$ and $\tau_2$ to obtain a match to the observed MDFs in Fig. 10. To carry out the numerical integration, we fix $M_0 = 1$ “unit” of initial gas mass and take $e = 0.05$ (that is, 5% of the available gas is converted to stars in each small timestep $\delta t$). The exact choice of $e$ is not critical; smaller fractions give a finer numerical mesh but maintain the same final MDF shape). After each timestep, the gas in the box is assumed as usual to be promptly mixed, so that all the stars formed in the next timestep then have the same abundance $Z(t)$ as the surrounding gas. After the integration is complete, the amount of stellar mass at each $Z$ is counted and the model MDF constructed.

### 4.2. Model Results and Comments

To match the steep observed rise of the MDF at low metallicity, we find that the amount of gas accreted in each timestep initially needs to be comparable with the amount converted to stars, i.e. $M_g(\text{added}) \sim e \cdot M_g$. After a handful of such steps, however, the infall rate must begin dying away to match the MDF peak and turnover region. The parameters for representative model solutions are summarized in Table 2, where $Z_g$ and $y$ are expressed in Solar units, $\tau_1$ and $\tau_2$ are expressed in numbers of timesteps, and the final galaxy mass $M_f$ is a multiple of the initial mass. Figures 15 and 16 show these sample results graphically. In all cases the final galaxy mass $M_f$ is 2 to 4 times bigger than the initial $M_0$. For the outer-halo fields, the effective yield must be about one-quarter Solar, while for the inner halo, it must be more than three times as high to generate the high-$Z$ section of the MDF.

For any $Z_g \lesssim 0.2 Z_\odot$ it is possible to find models which closely match the outer-halo MDF data.
However, for pre-enrichment values much above $Z_g \gtrsim 0.2Z_\odot$, plausible fits become more difficult to find, because the incoming gas is already too enriched to produce the correct numbers of low–$Z$ stars. We conclude, then, that within the context of an accretion model and in situ star formation, the incoming gas needs to have quite low metallicity.

The inner-halo MDF is more easily matched with slightly more enriched gas ($Z_g \gtrsim 0.1 - 0.2Z_\odot$), but none of these simple models can accurately reproduce the observed structure at $Z \gtrsim 0.7Z_\odot$. If the excess of high–$Z$ stars is real, then it may indicate that the metal-rich bulge component was created separately such as by a late merger, where the accreted material was already enriched.

The protogalactic clouds envisaged in our rough model correspond strikingly with the characteristics of the damped Ly\textalpha clouds. Recent high-resolution spectroscopic observations of these systems over a large range of redshifts (e.g., Pettini et al. 2000; Prochaska & Wolfe 1999, 2000) have made it increasingly clear that they are excellent candidates for protogalactic units out of which bigger galaxies assembled. The observed heavy-element abundances of these clouds cover a wide logarithmic range in [m/H] but are very metal-poor on average, with $\langle m/H \rangle \simeq -1.6$ (cf. the references cited above), corresponding to $\langle Z \rangle \sim 0.03Z_\odot$. This level would place them in a regime bracketed by our Figures 15 and 16.

The true mass infall rates cannot be derived directly from our model, because the “timestep” $\delta t$ is arbitrary and must be determined from external information. Unfortunately, the properties of the RGB stars which make them so useful as metallicity indicators also prevent them from supplying any very helpful information on their age distribution. Nevertheless, we can roughly calibrate the time sequence if we forcibly assume (following the hierarchical-merging simulations referred to above, as well as the evidence from the redshift distribution of the Ly\textalpha clouds) that the main bulk of star formation took place over a plausibly short interval such as $\sim 2$ Gy. For all the models, we find that it takes typically 50 – 80 timesteps with $e = 0.05$ to reach 80% of their final mass, suggesting $\langle \delta t \rangle \simeq 25 - 40$ Myr.

Finally, if the total stellar mass of NGC 5128 is $M \simeq 4.1 \times 10^{11}M_\odot$ (for an integrated magnitude $V_T = 6.2$ (van den Bergh 1976), a distance modulus $(m - M)_V = 28.3$, and $(M/L)_V = 7$), the amount of gas converted to stars in each timestep can be translated to a star formation rate (SFR) knowing $\delta t$. The corresponding average SFR in the first 2 Gy and integrated over the entire galaxy is $\simeq 150 M_\odot \, y^{-1}$. The maximum SFR for each model is listed at the end of Table 2. SFRs this large are an order of magnitude higher than observed in large galaxies at redshifts $z \sim 1$ (e.g., Postman et al. 2001), but are smaller than the rates observed in various extreme starburst systems such as the Ultra-Luminous Infrared Galaxies (Anantharamaiah et al. 2000; Dey et al. 1999; Scoville 2001; Alonso-Herrero et al. 1998).

---

4Strictly speaking, for the outer-halo fields we should define $M_f$ as the total mass of only the halo and not the entire galaxy, and similarly for the outer-bulge field we should use only the total mass of the bulge. However, the end result for the maximum SFR added up over the entire galaxy is the same in either case.
5. Globular Clusters and a New “Specific Frequency Problem”

As a final part of our discussion, we return to an issue first raised in Paper I: how does the metallicity distribution of the field-halo stars compare with that of the globular clusters in NGC 5128? Eventually, we would like to encompass the MDFs of both field-halo stars and clusters within a single formation history. To address this question, we take the samples of individually determined cluster metallicities published by G.Harris et al. (1992) (from the Washington $(C-T_1)$ index) and by Rejkuba (2001) (from the $(U-V)$ color index), each of which contain 60 to 70 clusters drawn from a wide range of radial distances in the halo. We view this as only a preliminary comparison, since considerably larger samples of clusters will soon be available from other wide-field imaging projects now in progress. Nevertheless, even with this rough comparison an interesting trend emerges.

5.1. Calibration of $(C-T_1)$ Metallicity Index

The Washington color $(C-T_1)$ is an effective metallicity index for globular clusters (Geisler & Forte 1990) and has been used as such in many studies. However, the original calibration of $(C-T_1)$ in terms of cluster metallicity employed Milky Way cluster reddenings and $[\text{Fe/H}]$ values which have been extensively revised and updated over the intervening years, to the point where it is of interest to redo the calibration. For the Milky Way cluster $(C-T_1)$ colors we use the fundamental list of H.Harris & Canterna (1977), while for their reddenings and modern $[\text{Fe/H}]$ values we use the 1999 edition of the W.Harris (1996) catalog, along with the reddening conversion $E(C-T_1) = 1.966E(B-V)$ (H.Harris & Canterna 1979). Combining this material for 48 clusters yields the relation shown in Figure 17.

With the contemporary data, the correlation between the two now appears to be mildly non-linear. A quadratic least squares solution of $(C-T_1)_0$ against $[\text{Fe/H}]$ (ignoring two obviously discrepant points at intermediate color, which belong to NGC 288 and 6522) yields

$$(C-T_1)_0 = 1.998 + 0.748[\text{Fe/H}] + 0.138[\text{Fe/H}]^2 \quad (3)$$

while the inverse solution gives

$$[\text{Fe/H}] = -6.037(1 - 0.82(C-T_1)_0 + 0.162(C-T_1)_0^2). \quad (4)$$

These reproduce the data with an rms scatter of ±0.16 dex in $[\text{Fe/H}]$, or ±0.055 mag in $(C-T_1)_0$. As Figure 17 shows, these two curves are almost identical except at the extreme low-metallicity end, which is not well constrained by the data and where in any case the color index becomes insensitive to metallicity. We use this new calibration to convert the $(C-T_1)$ indices of the NGC 5128 clusters (G.Harris et al. 1992) into metallicity.

For the Rejkuba (2001) sample, we use the correlation between $(U-V)_0$ and $[\text{Fe/H}]$ derived by Reed et al. (1994),

$$[\text{Fe/H}] = -3.061 + 2.015(U-V)_0. \quad (5)$$
The metallicity indices for the *globular clusters* are calibrated in terms of \([\text{Fe/H}]\), while our data for the *field stars* are in terms of \([\text{m/H}] = \log \left( \frac{Z}{Z_{\odot}} \right) \simeq [\text{Fe/H}] + [\alpha/\text{Fe}]\). Recent compilations of spectroscopic abundance measurements for old-halo giants in many galaxies (e.g. Shetrone et al. 2001) indicate that the \([\alpha/\text{Fe}]\) ratio has significant star-to-star scatter and that its mean value increases slowly with decreasing metallicity, staying in the range \(0.2 - 0.3\) for \(-3 \lesssim [\text{Fe/H}] \lesssim -0.5\). At more nearly Solar metallicity, spectra of Milky Way bulge stars and of the integrated bulge light of other large galaxies (McWilliam & Rich 1994; Trager et al. 2000) indicate that \([\alpha/\text{Fe}]\) remains near +0.2 though again with considerable object-to-object variance. Although it is not entirely clear whether this rather mixed ensemble of evidence should apply to NGC 5128 in particular, we adopt it as the best present guess. In summary, to convert our stellar data from \([\text{m/H}]\) to \([\text{Fe/H}]\) we adopt \([\alpha/\text{Fe}] = +0.2\) for \([\text{Fe/H}] \geq -0.6\), and then use a linear increase of \([\alpha/\text{Fe}]\) from +0.2 to +0.3 as \([\text{Fe/H}]\) decreases from −0.6 to −2.0.

### 5.2. Comparing Clusters With Field Stars

The available NGC 5128 globular cluster data include objects with projected galactocentric distances from \(R \simeq 2'\) to almost 22', and although the combined sample comprises only 127 clusters, it is already clear that the outer-halo objects are more metal-poor on average. More or less arbitrarily, we have divided the total set of clusters into two roughly equal subsets by radius (within \(R = 8'\) and beyond \(R = 8'\)) for comparison with the inner-halo and outer-halo field star samples, leaving about 65 clusters in each group.

The results are shown in Figure 18. The histograms for the clusters have been calculated with a smoothing kernel of 0.1 dex, while the stellar histograms have been smoothed by 0.05 dex, consistent with the internal uncertainties resulting from the random errors of photometry. In each graph, both curves have been normalized to the same total population of objects. The distinct lumpiness of the two cluster MDFs is a plain result of the still-small sample size and should not be ascribed any particular importance. For the inner halo, we find that the clusters and field stars have basically similar MDFs, with an extremely wide range of metallicities and broad peaks near \([\text{Fe/H}] \sim -0.5\). Given that the two distributions were established in somewhat different ways ((\(V-I\)) colors of RGB stars, versus \((C-T_1)\) integrated colors of clusters), we suggest that there is no strong reason to claim that they are fundamentally different.

For the outer halo, however, the clusters and field stars are strikingly different (in a formal statistical sense, they are different at the > 99.9% confidence level from a K-S test). The cluster MDF at least roughly resembles what we find for the globular clusters in the Milky Way and M31 (e.g. Barmby et al. 2000; W.Harris 2001), with a primary metal-poor peak at \([\text{Fe/H}] \simeq -1.6\) and a secondary group of more metal-rich clusters extending up to roughly Solar abundance. Unlike the Milky Way, however, the outer-halo stars lie at a considerably higher mean metallicity level, corresponding to about one order of magnitude in heavy-element abundance \(Z\). The case for resemblance to M31 is much stronger: as pointed out by Durrell et al. (2001) and W.Harris &
G.Harris (2001), the same dichotomy between the clusters and field-halo stars appears there, and to very much the same degree.

5.3. Specific Frequency: Discussion and Speculation

The ratio of numbers of globular clusters to field stars is defined as the specific frequency $S_N$ (W.Harris & van den Bergh 1981; W.Harris 2001). If it is reasonable to assume that the globular clusters at a given [Fe/H] can be associated with the field stars at the same metallicity, then our NGC 5128 data offer the chance to plot $S_N$ in a new way, as a function of the metallicity of the subpopulation. We did this roughly in Paper I by breaking the stars and clusters into just two metallicity groups, but the more comprehensive datasets now allow us to divide the metallicity range a bit further. Results for both the inner and outer halo, and for five metallicity bins in each zone of width $\Delta [\text{Fe/H}] = 0.4$ dex, are listed in Table 3 and plotted in Figure 19. In each bin, we integrate under the smoothed MDFs in Fig. 18 for the given metallicity range to estimate the relative numbers of stars and clusters. The resulting ratios are normalized so that the total $N_{cl}/N_\star$ over all metallicities equals the specific frequency $S_N(\text{total}) = 2.6$ for the whole galaxy (see Paper I). Evidence from wide-field starcounts (G.Harris et al. 1984) suggests that the globular cluster system follows the same overall radial distribution as the integrated halo light and thus the same global specific frequency can be used for both regions.

The actual zeropoints for each region are, however, less important than the trend with metallicity: we find that $S_N$ is a strong function of [Fe/H], changing by roughly a factor of three over the entire metallicity range. (The anomalously high $S_N$ for the most metal-rich clusters in the outer halo is likely to be strongly affected by small-sample statistics and is thus not very significant.) For [Fe/H] $\gtrsim -1.2$, we find $S_N$ values near 1–3, which are typical of those in field ellipticals, spirals, and recent merger products of disk galaxies (W.Harris 2001). For [Fe/H] $\lesssim -1.2$, the ratio is 4–8, comparable with many giant E galaxies, or (see below) many of the smallest dwarf ellipticals.

Metallicity-based differences in $S_N$ have been hinted at in previous work for other large galaxies such as M31 and NGC 4472 (see Durrell et al. 1994; Forbes et al. 1997; W.Harris 2001, as well as Paper I for NGC 5128 itself), but the existence of more extensive MDFs for both clusters and halo stars in the same system now allows us to quantify the trend much further. It suggests either that low-metallicity clusters formed at considerably higher efficiency than metal-richer ones, or that in some other way the formation of globular clusters did not go in lockstep with the field stars. This disparity is all the more puzzling, given the strong evidence for a global cluster formation efficiency of about 0.25% by mass that is remarkably similar in large galaxies of all types and sizes (McLaughlin 1999; Blakeslee 1999; Kavelaars 1999; W.Harris 2001).

For globular cluster systems there are at least three different kinds of “specific frequency problems”. The first of these, uncovered more than two decades ago, is the correlation of $S_N$ with E galaxy luminosity and the existence of central-supergiant galaxies (M87 and others like it) with
strikingly large cluster populations. A plausible solution to this issue (McLaughlin 1999; Blakeslee 1999; Kavelaars 1999) is that the $S_N$ values become much more nearly uniform when normalized to the total masses of the host galaxies including their X-ray halo gas. A second and almost equally old “specific frequency problem” is the overall large range in $S_N$ amongst rich cluster ellipticals, field ellipticals, and spirals (see W.Harris 2001, for a review). Understanding the origin of this range and variety is likely to involve numerous details of their formation histories including environmental factors such as the tendency of spiral/spiral mergers to form low-$S_N$ “field” ellipticals.

The correlation between $S_N$ and metallicity which we raise here is a third fundamental kind of “specific frequency problem”, and presents a new challenge to our understanding of the link between globular clusters and halo stars. Much observational and theoretical evidence indicates that (a) star clusters form within giant molecular clouds (GMCs), with perhaps a handful of clusters forming within any one GMC, and that (b) the typical masses of the protoclusters are about 3 orders of magnitude smaller than the total mass of the parent GMC itself (W.Harris & Pudritz 1994; McLaughlin & Pudritz 1996; McLaughlin 1999). The mean mass ratio $\epsilon_{cl} = 0.0025$ derived by McLaughlin (1999) corresponds to a baseline specific frequency $S_N \sim 5$ and is, in this hypothesis, the average mass ratio $M_{cl}/M_{GMC}$ of clusters formed within a typical GMC. Star clusters with masses $10^4 - 10^6 M_\odot$ (that is, globular clusters) should then form within GMCs of masses $10^7 - 10^9 M_\odot$. GMC masses in this range make them easily identifiable with protogalactic subsystems such as the Ly$\alpha$ clouds (Burgarella et al. 2001), the supergiant molecular clouds of W.Harris & Pudritz (1994), or the small cool-gas disks of Cole et al. (2000) which merge hierarchically to build larger galaxies at the same time as they are forming their own embedded star clusters.

As did Forbes et al. (1997), W.Harris (2001), and Burgarella et al. (2001), we imagine a sequence of events in which the first major round of globular cluster and star formation took place within these dwarf-galaxy-sized gas clouds, scattered throughout the much larger potential well of the giant proto-elliptical. These first clusters and field stars took on the low metallicities of their primordial host clouds. As the clouds merged hierarchically, later continuous rounds of star formation took place in which the progressively more enriched gas formed the metal-richer clusters and field stars. But some additional feature of the story is needed to force the ratio of field stars to globular clusters steadily upward (or conversely, to push $S_N$ steadily downward) in these later rounds of star formation, as the evidence of Fig. 19 would indicate.

We suggest that this additional feature may be the relative timing of globular cluster formation versus their associated field stars. If the clusters, which should form out of the densest, most massive, or highest-pressure clumps of gas within their host GMCs, are constructed $soonest$ in the sequence – that is, ahead of the majority of the field stars – then the effective $S_N$ emerging from the entire GMC will depend on when the star formation history is truncated. Earlier versions of this idea were suggested by Durrell et al. (1996) and W.Harris (2001).

The essential concept is illustrated schematically in Figure 20. Within a given protogalactic cloud (supergiant GMC), the clusters are envisaged to form predominantly near the beginning of
the GMC’s history. By contrast, the general star formation distributed everywhere throughout the cloud in less dense local regions and at lower conversion efficiency is assumed to take longer to ramp up and then die away. A “normal” specific frequency (number of clusters per unit field-star mass) would result if the host GMC is able to complete its full sequence of star formation (in Fig. 20, up to time $T_1$). However, if the normal sequence is truncated at an earlier time $T_2$ – for example, if the GMC itself is disrupted and its remaining gas ejected – then the net specific frequency can be considerably larger: the same number of globular clusters is still present, having formed very early on, but most of the general stellar population does not get a chance to form.

This premature truncation of the internal enrichment and star formation is exactly what would be expected for isolated protogalactic clouds of masses $\lesssim 10^9 M_\odot$: as first calculated by Dekel & Silk (1986), the initial major round of supernovae and stellar winds in such dwarf-galaxy-sized systems would be sufficient to drive the rest of the gas out of their modest potential wells and terminate any further star formation. A prediction emerging from this combined picture is that dwarf ellipticals, which can plausibly be viewed to have evolved from single protogalactic clouds in isolation, should have progressively larger specific frequencies for smaller dwarfs (in addition, of course, to the long-established observational trend that smaller dwarfs should have lower mean metallicity). The available $S_N$ data for dE galaxies (Durrell et al. 1996; Miller et al. 1998) are strongly consistent with this prediction, as shown by McLaughlin (1999). The smallest dE,N systems have $S_N$ reaching the remarkably high range $\sim 15 – 25$ (though with considerable scatter), while the largest dwarfs have $S_N$ more consistently at “normal” levels of $\sim 3 – 6$ (see the discussions of McLaughlin 1999; W.Harris 2001). Another piece of observational evidence that is, perhaps, related to this picture is that very high specific frequencies or specific luminosities have been derived for some starburst galaxies (e.g., Meurer et al. 1995; Zepf et al. 1999; Larsen & Richtler 2000). The cluster-to-field-star ratio very early in the burst will look nominally larger than normal if the clusters form preferentially earlier. Nevertheless, it is intriguing that most such systems (see Larsen & Richtler 2000, for a compilation) have specific luminosities at the 1% level or less, comparable with the “universal” 0.25% efficiency ratio for old cluster systems.

We suggest that the high $S_N$ levels exhibited at low metallicities in giant galaxies (Fig. 19 and the discussion above) may be the result of the same basic phenomenon. At early times, the protogalactic dwarfs were dispersed widely enough that they could, for a short time, evolve almost in isolation. They formed metal-poor clusters and some stars, but then ejected most of their gas, leaving behind a low-metallicity but high–$S_N$ stellar population. The large amount of leftover gas then re-collected in the much bigger potential well of the giant galaxy as hierarchical merging continued, so that in the later rounds of star formation most of the gas could be kept in situ and the star formation could run to completion, yielding a normal ratio of clusters to field stars.

In this admittedly speculative picture, the overall level of heavy-element enrichment is still closely connected with chronology within any one GMC. If the later rounds of star formation behave the same way as the first round (that is, if globular clusters form soonest within any GMC regardless of its initial metallicity), then we would expect that the highest-metallicity field stars in
the bulge of the galaxy should be more metal-rich than the highest-metallicity globular clusters. In other words, the MDF for the field-star population should extend to a higher “top end” than that of the clusters. In NGC 5128, it should be possible to test this prediction with a more comprehensive set of cluster data, along with efforts to extend the field-star MDF to higher metallicity limits than we were able to do in the present study.

If the timing argument we have outlined above proves to be wrong, then the obvious alternative is to invoke an intrinsic dependence of cluster formation efficiency on metallicity. That is, we would in that case assume that the very low−Z protogalactic clouds inhabiting the early universe converted their gas into dense, bound star clusters about 3 times more efficiently than did the later generations of GMCs which had been enriched to levels higher than about one-tenth Solar abundance. A third possibility, recently revived by Cen (2001), would be to assume arbitrarily that the lowest-metallicity clusters were formed in the early pregalactic universe by way of a fundamentally different route (in Cen’s model, at the epoch of cosmological reionization). This latter route appears to us to be much less attractive since it does not relate to any currently visible evidence for star and cluster formation in GMCs, nor does it explain the otherwise-similar properties of globular clusters (mass distribution functions, structural parameters) at all metallicities and in all types of galaxies.

6. Summary

We present new HST/WFPC2 photometry in \((V, I)\) for a starfield 8 kpc from the center of the nearby giant E galaxy NGC 5128. By using a grid of red-giant evolutionary tracks finely spaced in metallicity, we convert the color-magnitude diagram of the stellar population into a metallicity distribution function, and combine it with the MDFs for our previously analyzed outer-halo fields at 21 and 31 kpc distance (Papers I and II).

In all three fields, the CMD is completely dominated by a conventionally “old” (many Gy) stellar population of red giants, with moderately high mean metallicity \(\langle m/H \rangle \sim -0.5\) and very few low-metallicity stars like the ones dominating the Milky Way halo. However, the innermost of our three fields is distinctly more metal-rich than the outer ones; if we arbitrarily view it as consisting of two sub-populations (a halo component like the outer fields, plus a bulge component), then the “bulge” by itself has a mean metallicity \([m/H] \sim -0.2\), similar to bulge compositions in many other large elliptical and spiral galaxies.

We find that a simple accreting-box chemical evolution model that assumes a gas infall rate that is initially rapid but then declines exponentially with time is capable of closely matching the MDF of the outer-halo fields, as long as the infall gas has low metallicity \([m/H] \sim -1\) or less) and the effective nucleosynthetic yield is \(y \sim 0.3Z_\odot\). The same type of model with \(y \sim 0.8Z_\odot\) can also match the inner-halo field.

Finally, we compare our field-star MDFs to the MDF for the old globular clusters in NGC 5128,
with both samples divided into inner and outer regions. In both regions, there is a striking increase of net specific frequency (number of clusters per unit halo field-star numbers) at low [Fe/H], with the lowest-metallicity bins having 3 times more clusters per unit halo light than the highest-metallicity bins. We suggest that a way to explain this trend, within the context of a hierarchical-merging formation picture, is that globular clusters form preferentially sooner within any one starburst than the accompanying field stars. The first clusters to form were in relatively isolated, small, low–Z protogalactic clouds, which ejected most of their gas after the first round of star formation, leaving behind a population of low-metallicity clusters and only a few field stars. In the later rounds of star formation after the merging process had advanced much further, the sites of star and cluster formation were in the much deeper potential well of the emerging giant elliptical. They could thus hold onto much higher fractions of their gas, allowing star formation to run to completion and producing proportionally larger numbers of high-metallicity stars for each globular cluster.

This work was supported by the Natural Sciences and Engineering Research Council of Canada through research grants to the authors. We are pleased to acknowledge the hospitality and support at Mount Stromlo Observatory (RSAA/ANU) during the authors’ research leaves when this paper was written.

REFERENCES

Alonso-Herrero, A., Rieke, M.J., & Rieke, G.H. 1998, Ap.SpaceSci, 263, 131
Anantharamaiah, K.R., Viallefond, F., Mohan, N.R., Goss, W.M., & Zhao, J.H. 2000, ApJ, 537, 613
Ashman, K.M., & Zepf, S.E. 1992, ApJ, 384, 50
Barbuy, B., Renzini, A., Ortolani, S, Bica, E., & Guarnieri, M.D. 1999, A&A, 341, 539
Barmby, P., Huchra, J.P., Brodie, J.P., Forbes, D.A., Schroder, L.L., & Grillmair, C.J. 2000, AJ, 119, 727
Barnes, J.E., & Hernquist, L. 1992, ARA&A, 30, 705
Beasley, M., Harris, W.E., Harris, G.L.H., & Forbes, D. 2002, in preparation
Beaulieu, S.F. et al. 2001, AJ, 121, 2618
Bertelli, G., Bressan, A., Chiosi, C., Fagotto, F., & Nasi, E. 1994, A&AS, 106, 275
Binney, J., & Merrifield, M. 1998, Galactic Astronomy (Princeton University Press)
Blakeslee, J. 1999, AJ, 118, 1506
Burgarella, D., Kissler-Patig, M., & Buat, V. 2001, AJ, 121, 2647
Carretta, E., Cohen, J.G., Gratton, R.G., & Behr, B.B. 2001, AJ, 122, 1469
Cen, R. 2001, ApJ, 560, 592
Chaboyer, B., Green, E.M., & Liebert, J. 1999, AJ, 117, 1360
Chiappini, C., Matteucci, F., & Gratton, R. 1997, ApJ, 477, 765
Cohen, J.G., Gratton, R.G., Behr, B.B., & Carretta, E. 1999, ApJ, 523, 739
Cole, S., Lacey, C.G., Baugh, C.M., & Frenk, C.S. 2000, MNRAS, 319, 168
Côté, P., Marzke, R.O., & West, M.J. 1998, ApJ, 501, 554
Côté, P., Marzke, R.O., West, M.J., & Minniti, D. 2000, ApJ, 533, 869
Côté, P., West, M.J., & Marzke, R.O. 2001, ApJ, in press
Dekel, A., & Silk, J. 1986, ApJ, 303, 39
Dey, A., Graham, J.A., Ivison, R.J., Smail, I., Wright, G.S., & Liu, M.C. 1999, ApJ, 519, 610
Dubinski, J. 1998, ApJ, 502, 141
Durrell, P.R., Harris, W.E., & Pritchet, C.J. 1994, AJ, 108, 2114
Durrell, P.R., Harris, W.E., Geisler, D., & Pudritz, R.E. 1996, AJ, 112, 972
Durrell, P.R., Harris, W.E., & Pritchet, C.J. 2001, AJ, 121, 2557
Fleming, D.E.B., Harris, W.E., Pritchet, C.J., & Hanes, D.A. 1995, AJ, 109, 1044
Forbes, D.A., Brodie, J.P., & Grillmair, C.J. 1997, AJ, 113, 1625
Frogel, J. A., Tiede, G. P., & Kuchinski, L. E. 1999, AJ, 117, 2296
Geisler, D., & Forte, J.C. 1990, ApJ, 350, L5
Harris, G.L.H., Geisler, D., Harris, H.C., & Hesser, J.E. 1992, AJ, 104, 613
Harris, G.L.H., Harris, W.E., & Poole, G. B. 1999, AJ, 117, 855 (Paper I)
Harris, G.L.H., & Harris, W.E. 2000, AJ, 120, 2423 (Paper II)
Harris, G.L.H., Hesser, J.E., Harris, H.C., & Curry, P.J. 1984, ApJ, 287, 175
Harris, H.C., & Canterna, R. 1977, AJ, 82, 798
Harris, H.C., & Canterna, R. 1979, AJ, 84, 1750
Harris, W.E. 1996, AJ, 112, 1487
Harris, W.E. 2001, in Star Clusters, Saas-Fee Advanced Course 28 (New York: Springer), ed. L.Labhardt & B.Binggeli.
Harris, W.E., & Harris, G.L.H. 2001, AJ, 122, in press
Harris, W.E., & Pudritz, R.E. 1994, ApJ, 429, 177
Harris, W.E., & van den Bergh, S. 1981, AJ, 86, 1627
Hibbard, J.E., & Yun, M.S. 1999, ApJ, 522, L93
Holtzman, J. A. et al. 1995, PASP, 107, 1065
Ibata, R. A., & Gilmore, G. F. 1995, MNRAS, 275, 605
Kavelaars, J.J. 1999, in Galaxy Dynamics, ASP Conference Series 182, ed. D.Merritt, M.Valluri, & J.A.Sellwood (San Francisco: ASP), 437
Kauffmann, G., White, S. D. M., & Guideroni, B. 1993, MNRAS, 264, 201
Lacey, C.G., & Fall, S.M. 1983, MNRAS, 204, 791
Larsen, S.S., & Richtler, T. 2000, A&A, 354, 836
Larson, R.B. 1972, Nature Phys.Sci., 236, 7
Marleau, F. R., Graham, J. R., Liu, M. C., & Charlot, S. 2000, AJ, 120, 1779
McLaughlin, D.E. 1999, AJ, 117, 2398
McLaughlin, D.E., & Pudritz, R.E. 1996, ApJ, 457, 578
McWilliam, A., & Rich, R. M. 1994, ApJS, 91, 749
Meurer, G.R., Heckman, T.M., Leitherer, C., Kinney, A., Robert, C., & Garnett, D.R. 1995, AJ, 110, 2665
Miller, B.W., Lotz, J.M., Ferguson, H.C., Stiavelli, M., & Whitmore, B.C. 1998, ApJ, 508, L133
Naab, T., & Burkert, A. 2001, ApJ, 555, L91
Origlia, L., Rich, R.M., & Castro, S.M. 2001, Bull.AAS, 199, 56.16
Pagel, B. E. J., & Patchett, B. E. 1975, MNRAS, 172, 13
Pearce, F. R. 1999, ApJ, 521, L99
Peng, E. 2001, private communication
Peterson, R.C., & Green, E.M. 1998, ApJ, 502, L39
Pettini, M., Ellison, S.L., Steidel, C.C., Shapley, A.E., & Bowen, D.V. 2000, ApJ, 532, 65
Portinari, L., Chiosi, C., & Bressan, A. 1998, A&A, 334, 50
Postman, M., Lubin, L.M., & Oke, J.B. 2001, AJ, 122, 1125
Prochaska, J.X., & Wolfe, A.M. 1999, ApJS, 121, 369
Prochaska, J.X., & Wolfe, A.M. 2000, ApJ, 533, L5
Reed, L.G., Harris, G.L.H., & Harris, W.E. 1994, AJ, 107, 555
Rejkuba, M. 2001, A&A, 369, 812
Salaris, M., & Weiss, A. 1998, A&A, 335, 943
Salasnich, B., Girardi, L., Weiss, A., & Chiosi, C. 2000, A&A, 361, 1023
Schweizer, F. 1987, in Nearly Normal Galaxies, Proc. 8th Santa Cruz Summer Workshop in Astronomy and Astrophysics (New York: Springer-Verlag), p. 18
Scoville, N. 2001, BAAS, 198.3410
Shetrone, M.D., Côté, P., & Sargent, W.L.W. 2001, ApJ, 548, 592
Soria, R. et al. 1996, ApJ, 465, 79
Stetson, P.B. 1992, in Astronomical Data Analysis Software and Systems I, ASP Conf.Ser. Vol.8, edited by G.H.Jacoby (ASP, San Francisco), p. 289
Taylor, B.J. 2001, A&A, 377, 473
Theis, Ch., Burkert, A., & Hensler, G. 1992, A&A, 265, 465
Trager, S.C., Faber, S.M., Worthey, G., & Gonzalez, J.J. 2000, AJ, 119, 1645
VandenBerg, D. A., Swenson, F. J., Rogers, F. J., Iglesias, C. A., & Alexander, D. R. 2000, ApJ, 532, 430
van den Bergh, S. 1976, ApJ, 208, 673
Yi, S., Demarque, P., Kim, Y.-C., Lee, Y.-W., Ree, C.-H., Lejeune, Th., & Barnes, S. 2001, ApJS, 136, 417
Zepf, S.E., Ashman, K.M., English, J., Freeman, K.C., & Sharples, R.M. 1999, AJ, 118, 752
Zepf, S.E. et al. 2000, AJ, 120, 2928

This preprint was prepared with the AAS L\LaTeX macros v5.0.
| [m/H] | n(Outer) | n(Inner) | [m/H] | n(Outer) | n(Inner) |
|-------|----------|----------|-------|----------|----------|
| -2.65 | 2.1      | 2.1      | -1.05 | 210.7    | 113.1    |
| -2.55 | 7.3      | 2.1      | -0.95 | 282.4    | 135.7    |
| -2.45 | 3.2      | 5.2      | -0.85 | 309.1    | 175.4    |
| -2.35 | 5.2      | 2.1      | -0.75 | 407.9    | 235.0    |
| -2.25 | 15.7     | 3.1      | -0.65 | 530.9    | 320.3    |
| -2.15 | 14.6     | 6.2      | -0.55 | 549.5    | 428.7    |
| -2.05 | 18.8     | 11.5     | -0.45 | 614.2    | 497.7    |
| -1.95 | 27.1     | 10.4     | -0.35 | 467.7    | 495.4    |
| -1.85 | 37.6     | 15.7     | -0.25 | 332.1    | 464.6    |
| -1.75 | 40.7     | 16.7     | -0.15 | 148.3    | 429.5    |
| -1.65 | 42.0     | 23.1     | -0.05 | 25.4     | 304.2    |
| -1.55 | 49.2     | 23.0     | 0.05  | 1.9      | 340.6    |
| -1.45 | 74.2     | 38.7     | 0.15  | 0.0      | 160.6    |
| -1.35 | 87.8     | 23.0     | 0.25  | 0.0      | 72.0     |
| -1.25 | 111.7    | 56.5     | 0.35  | 0.0      | 40.0     |
| -1.15 | 170.7    | 81.7     | 0.45  | 0.0      | 4.0      |
Table 2. Fitting Parameters for Accreting Box Models

| Parameter           | Outer Fields | Inner Field | Outer Fields | Inner Field |
|---------------------|--------------|-------------|--------------|-------------|
| $Z_g/Z_\odot$       | 0.0          | 0.0         | 0.2          | 0.2         |
| $y/Z_\odot$         | 0.32         | 0.87        | 0.25         | 0.81        |
| $\tau_1/\delta t$  | 7            | 5           | 1            | 0           |
| $\tau_2/\delta t$  | 35           | 20          | 19           | 12          |
| $M_f/M_0$           | 3.5          | 3.9         | 1.9          | 2.4         |
| max SFR ($M_\odot/y$) | 225          | 155         | 240          | 158         |

Table 3. Specific Frequency vs. Metallicity

| [Fe/H] Range | $S_N$ (inner halo) | $S_N$ (outer halo) |
|--------------|--------------------|--------------------|
| $<-1.6$      | $3.4 \pm 1.2$      | $8.3 \pm 1.9$      |
| $(-1.6, -1.2)$ | $4.4 \pm 1.5$     | $4.9 \pm 1.2$      |
| $(-1.2, -0.8)$ | $4.0 \pm 1.1$     | $1.3 \pm 0.4$      |
| $(-0.8, -0.4)$ | $1.6 \pm 0.4$     | $0.9 \pm 0.3$      |
| $>-0.4$      | $1.6 \pm 0.4$      | $10.3 \pm 3.1$     |
Fig. 1.— Color-magnitude diagram for 17,326 measured stars in the PC1 field, centered a projected distance of 8 kpc southwest of the center of NGC 5128. The dashed line indicates the magnitude limit of 50% detection completeness (see text for explanation).

Fig. 2.— Artificial-star tests for the DAOPHOT/ALLSTAR photometry in the PC1 field. Stars are added 500 at a time to the original image with \((I, V - I)\) values along the dispersionless line (left panel). Their measured values (right panel) scatter symmetrically about the input relation. The dashed line shows the 50% detection completeness boundary.

Fig. 3.— The two panels show, for the artificial-star tests, the scatter of measured magnitudes versus input magnitude (\(\Delta I\) versus \(I\), \(\Delta V\) versus \(V\)). Although the random uncertainty increases with magnitude, no systematic bias is present over the upper magnitude range used for the definition of the metallicity distribution function (see text).

Fig. 4.— Detection completeness of the photometry as a function of magnitude, derived from the artificial-star tests. Here the completeness fraction \(f\) is defined as the number of measured stars in a given magnitude bin relative to the number of input stars in that bin. The two smooth curves are Pritchet interpolation functions (Fleming et al. 1995). The curves shown are for mean colors \((V - I) = 2\) and are only to be taken as indicative of the average magnitudes at which the photometry becomes severely incomplete.

Fig. 5.— Calibration grid for the derivation of stellar metallicities. The open circles indicate fiducial points for the seven “standard” Milky Way clusters described in the text (the three most metal-rich, NGC 104, 6553, and 6791 are labelled). The solid lines are the evolutionary tracks for the \(\alpha\)-enhanced models of VandenBerg et al. (2000) used in Paper II, while the two dashed lines are estimated tracks for heavy-element abundances \(Z = 0.025\) and 0.045 as described in the text.
Fig. 6.— Color-magnitude array for the 8 kpc field, plotted in terms of bolometric magnitude versus intrinsic color. The RGB model grid from VandenBerg et al. (2000; solid lines) are superimposed on the data. The two dashed lines are two more metal-rich tracks defined as described in the text. Interpolation within the grid is used to estimate the metallicity of each star.

Fig. 7.— Metallicity distribution function (MDF) for the combined outer-halo fields (21 kpc and 31 kpc) discussed in Papers I and II. The results have been recalculated from the stellar model grid of Figure 5 through the updated interpolation code as described in the text. The MDFs for four different luminosity intervals along the RGB are shown, from the top of the RGB down to a point two magnitudes lower. Note the increased spread at fainter levels, due to increased photometric scatter. The unshaded histograms show the completeness-corrected results as described in the text.

Fig. 8.— MDF for the inner-halo 8 kpc field, calculated in the same way as for the outer-halo fields in the previous figure. The shaded histograms represent the raw number counts, while the higher unshaded histogram line represents the totals corrected for detection incompleteness as described in the text.

Fig. 9.— Final MDFs for the outer-halo fields (upper panel) and the inner-halo field (lower panel). Table 1 contains the same data, tabulated in 0.1 dex intervals. All RGB stars brighter than $M_{bol} = -2.5$, i.e. the upper 1.5 magnitudes of the RGB, are used here in an unweighted sum. In each panel the shaded histogram is corrected for photometric incompleteness, while the solid line underneath shows the raw, uncorrected total.

Fig. 10.— Heavy-element abundance distribution for the inner-halo (solid line) and outer-halo (dashed line) fields, presented as a probability distribution as defined in the text. Here, $p(Z)dZ$ is the probability of finding a star with metallicity in the range $(Z, Z + dZ)$ and the integral over all $Z$ is normalized to 1.000. In this particular scaling, the interval $dZ$ is equal to 0.01$Z_{\odot}$. The dotted line at bottom shows the distribution for the inner field without any correction for photometric incompleteness.
Fig. 11.— Normalization of the metal-poor section of the MDF in the inner and outer fields. Here $n(\text{inner})/n(\text{outer})$ is the ratio of the numbers of stars in each field with [$m/H$] less than the given cutoff value. For [$m/H$] $\lesssim -1$ the ratio is nearly constant (see text).

Fig. 12.— Upper panel: The MDFs for the inner (unshaded histogram) and outer (shaded histogram) regions of the NGC 5128 halo. The outer-halo MDF has been normalized to match the inner-halo one for the metal-poor range [$m/H$] $< -1$ defined in Figure 11. Lower panel: The residual MDF after subtraction of the scaled outer-halo region from the inner-halo one. This differential MDF by hypothesis represents the outer bulge stellar population of NGC 5128. The roughly normal distribution of the residual MDF is indicated by the Gaussian curve with mean at [$m/H$] = -0.19 and standard deviation 0.25 dex.

Fig. 13.— The “delayed exponential” gas infall for the accreting-box model used in the text. Gas is assumed to fall in at a constant rate $k$ for a time $\tau_1$, then its infall rate declines exponentially with an e-folding time $\tau_2$.

Fig. 14.— A schematic MDF resulting from an accreting-box model (solid line). The parameters $k$ and $\tau_1$ determine the height and extent of the initial rise at lowest metallicity; $\tau_2$ determines the shape of the turnover region at the peak; and the yield rate $y$ determines the steepness of the decline at high metallicity. The closed-box model (dashed line) corresponds to $k = 0$, and follows an exponential decline from the start.

Fig. 15.— Sample fits of the accreting-box chemical evolution model to the inner- and outer-halo MDFs. We use here the probability distribution form of the MDF in terms of linear abundance $Z$. Solid lines show the observed MDFs from Figure 10, while the dashed lines show the models with parameters as listed in Table 2. These models assume the accreted gas has “primordial” abundance, $Z_g = 0$. 
Fig. 16.— Same as the previous figure, but for accreted gas with abundance $Z_g = 0.2Z_\odot ([\text{m/H}] = -0.7)$.

Fig. 17.— Calibration of the $(C - T_1)_0$ integrated color against metallicity, for 48 Milky Way clusters. The solid line is the quadratic relation from Equation 5 given in the text, the dashed line is from Equation 6, and the short-dotted line is the original linear relation of Geisler & Forte (1990).

Fig. 18.— Comparison of the metallicity distributions in [Fe/H] for the field halo stars (dashed lines) and NGC 5128 globular clusters (solid lines). Each distribution has been smoothed as described in the text. The upper panel shows the 8-kpc halo field compared with the globular clusters at projected distances less than 8 arcmin from the center of the galaxy. The lower panel shows the two outer-halo fields (21 kpc and 31 kpc) compared with the globular clusters outward of 8 arcmin.

Fig. 19.— Specific frequency $S_N$ of the globular cluster population in NGC 5128, plotted as a function of metallicity interval [Fe/H]. The data points are listed in Table 3, separately for the inner-halo ($R < 8'$) and outer-halo ($R > 8'$) regions. The dashed line at $S_N = 2.6$ is the global average for the entire galaxy.

Fig. 20.— Schematic concept for star and cluster formation within a single GMC or pregalactic gas cloud. (In reality the formation rates would be much more segmented and “burst”-driven than the smooth curves shown here.) If the gas supply in the cloud continues star formation till time $T_1$ (dashed line at right) a normal ratio $S_N$ of clusters to field stars will result. However, if the formation is truncated at some earlier time $T_2$, then the cloud will end up with a high effective specific frequency.
This figure "Harris.fig1.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig2.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig3.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig4.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig5.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig6.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig7.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig8.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig9.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig10.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig11.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig12.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig13.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig14.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig15.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig16.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig17.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig18.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig19.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1
This figure "Harris.fig20.jpg" is available in "jpg" format from:

http://arxiv.org/ps/astro-ph/0204440v1