MEASURING THE ANGULAR CORRELATION FUNCTION FOR FAINT GALAXIES IN HIGH GALACTIC LATITUDE FIELDS

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ABSTRACT

A photometric survey of faint galaxies in three high Galactic latitude fields (each \( \sim 49 \text{ arcmin}^2 \)) with subarcsecond seeing is used to study the clustering properties of the faint galaxy population. Multicolor photometry of the galaxies has been obtained to magnitude limits of \( V \sim 25, R \sim 25, \) and \( I \sim 24 \). Angular correlation analysis is applied to magnitude-limited and color-selected samples of galaxies from the three fields for angular separations ranging from \( 10'' \) to \( 126'' \). General agreement is obtained with other recent studies, which show that the amplitude of the angular correlation function, \( \omega(\theta) \), is smoothly decreasing as a function of limiting magnitude. The observed decline of \( \omega(\theta) \) rules out the viability of "maximal merger" galaxy evolution models.

Using redshift distributions extrapolated to faint magnitude limits, models of galaxy clustering evolution are calculated and compared to the observed \( J \) band \( \omega(\theta) \). Faint galaxies are determined to have correlation lengths and clustering evolution parameters of either \( r_0 = 4 h^{-1} \) Mpc and \( e \sim 0-1 \); \( r_0 = 5-6 h^{-1} \) Mpc and \( e > 1 \); or \( r_0 = 2-3 h^{-1} \) Mpc and \( e \sim -1.2 \), assuming \( q_0 = 0.5 \) and with \( h = H_0/100 \) km s\(^{-1}\) Mpc\(^{-1}\). The latter case is for clustering fixed in comoving coordinates and is probably unrealistic since most local galaxies are observed to be more strongly clustered. Even though the first of the three cases has the most reasonable rate of clustering evolution, distinguishing the correct \( r_0 \) for the faint galaxies is not possible with the current data. No significant variations in the clustering amplitude as a function of color are detected, for all the color-selected galaxy samples considered. The validity of this result is discussed in relation to other determinations of \( \omega(\theta) \) for galaxies selected by color.

Subject headings: cosmology: large-scale structure of universe — cosmology: observations — galaxies: clusters: general — galaxies: evolution — galaxies: photometry

1. INTRODUCTION

Angular correlation function analysis of large photometric samples of galaxies is a popular tool for quantifying the large-scale structure of the local universe. Its power as a diagnostic of the galaxy distribution lies in the simplicity of its application, essentially requiring only a counting of pairs of galaxies at given angular separations and normalizing these results with respect to the number of pairs expected for a random distribution. The angular correlation function, \( \omega(\theta) \), is a two-dimensional analogue of the spatial two-point correlation function, \( \xi(r) \), where the latter quantity has the important property of being the Fourier transform of the power spectrum, \( P(k) \), of the galaxy distribution. Measurements of either \( \xi(r) \) or \( P(k) \) are important tests of structure formation models, such as cold dark matter (CDM) scenarios (Davis et al. 1985) and are also necessary for understanding the relationship between nearby, bright galaxies and their faint, distant counterparts, along with the evolutionary processes that link the two populations.

The angular and spatial two-point correlation functions for bright magnitude-limited samples of nearby galaxies have been determined, in various forms, by several authors over the last few decades (see Peebles 1980, 1993, and references therein). Numerous studies have measured \( \xi(r) \) to be a power-law, \( (r/r_0)^{-\gamma} \), with \( \gamma \sim 1.7-1.8 \) and the correlation length, \( r_0(z = 0) \), is estimated to be \( \sim 5 h^{-1} \) Mpc for local galaxy populations (Davis & Peebles 1983; Loveday et al. 1992). There is still considerable uncertainty concerning the exact normalization and slope, owing to possible systematic errors from morphological mixing, redshift distortions from clusters of galaxies and differing approaches to the correlation analysis (Bernstein et al. 1994; Loveday et al. 1995).

Another potential problem for studies of nearby galaxies may be that a significant population of low surface brightness galaxies is being systematically ignored, due to high surface brightness selection effects, in contrast to faint galaxy surveys where this is not a problem (McGaugh 1994).

Pioneering photographic studies of faint galaxies done by Phillipps et al. (1978), Koo & Szalay (1984), Stevenson et al. (1985), and Pritchet & Infante (1986) determined \( \omega(\theta) \) down to magnitude limits of \( B < 23-24 \). Only in the last decade has it been possible to measure \( \omega(\theta) \) for fainter galaxy samples because of the emergence of CCD imaging cameras, in particular the large-format devices. Efstathiou et al. (1991), hereafter EBKTG, found the faint blue galaxy population in their CCD images (for \( 24 < B_J < 26 \) ) to be weakly clustered at \( 30'' \) separations relative to local galaxy populations. They concluded that either (1) most of the faint galaxies were members of an hitherto unobserved population that had faded away by the current epoch, (2) galaxy clustering was insufficiently described by basic models of gravitational instability or (3) that spacetime geometry departed significantly from an Einstein–de Sitter universe. However, as Koo & Kron (1992) point out, the EBKTG result implicitly assumes that galaxies with different morphologies have similar intrinsic clustering properties. This is not the case locally (Loveday et al. 1995; Giovanelli, Haynes, & Chincarini 1986; Davis & Geller 1976), and clearly this is an effect that needs to be addressed by using

\(^{1}\) Visiting Astronomer, Canada-France-Hawaii Telescope (CFHT), operated by the National Research Council of Canada, le Centre National de la Recherche Scientifique de France, and the University of Hawaii.
various sample selection criteria for faint galaxies such as morphologies and colors, in addition to magnitude limits.

Neuschaefer, Windhorst, & Dressler (1991, hereafter NWD) measured $\omega(\theta)$ down to $g \sim 25$ and found a similar amplitude as EBKTG at $g \sim 24.8$. The monotonic decrease of the amplitude of $\omega(\theta)$ for a given angular separation as a function of survey magnitude limit demonstrated by NWD has been observed in a number of other studies (Pritchet function of survey magnitude limit demonstrated by NWD of the amplitude of $\omega(\theta)$ as at $g \sim 24.8$. The monotonic decrease of the amplitude of the angular correlation function for the extremely blue and red subsets of galaxies with $20 < B_j < 23.5$ in their sample of 5900 galaxies. LSK claim the $\omega(\theta)$ excess for the reddest galaxies is due to intrinsic clustering of the sample (morphology-density relation) since most of these objects are probably E/S0 galaxies and both fields contain known clusters. The increase of the clustering amplitude for the bluest objects in the sample is explained as being caused by a faint population of galaxies with $z < 0.3$. Roche et al. (1996) calculated $\omega(\theta)$ for color-selected samples to significantly fainter magnitude limits ($B \sim 25.5$ and $R \sim 24.5$) than LSK. For $\sim 7000$ galaxies, they determined the amplitude of $\omega(\theta)$ for the red $[(B-R) > 1.5]$ sample to be higher than that calculated for the blue $[(B-R) < 1.5]$ galaxies. This result led Roche et al. to suggest that the decrease in the amplitude of $\omega(\theta)$ for all galaxies with $B > 23$ is caused by the same blue galaxies that are responsible for the number counts exceed around this magnitude range. Using a pure luminosity evolution model they explain the correlation function color dependence at $B \sim 25$ as being due to red galaxies with $z < 1$, in addition to the blue galaxy sample consisting of both late-type dwarfs at low/moderate redshifts and evolving $L^*$ galaxies having redshifts from $z \sim 0.5$–3. In this paper, $\omega(\theta)$ is calculated for a sample of galaxies imaged in $V$, $R$, and $I$ to respective magnitude limits of 25, 25, and 24, combining data from three different high Galactic latitude fields. Woods, Fahlman, & Richer (1995) determined the close pair fraction for one of the fields (NF1) studied here. This nearest neighbor approach is complementary to the correlation analysis since it measures clustering behavior at the smallest possible angular separations, while the greater numbers of galaxies analyzed in this paper can be used to estimate the angular correlation function over a range of larger angular separations (as an aside, note that eqs. [1] and [2] in Woods et al. 1995 are both missing a factor of $\rho$ in the integrand). Galaxy samples presented here are among the deepest yet used for measurements of $\omega(\theta)$. Also, the multibandpass data allow the clustering variations with color to be measured for these faint galaxies. Observations and preliminary data reduction and analysis techniques are outlined in §§ 2 and 3, respectively. The adopted approach for estimating the angular correlation function, along with a summary of the galaxy clustering model used, is presented in § 4. Clustering results for magnitude-limited and color-selected samples of galaxies are given in § 5. This section also contains a comparison of the clustering detected in the $I$ band to models of the spatial correlation function, which are calculated with extrapolated redshift distributions provided by the CFRS (Lilly et al. 1995a). Finally, possible interpretations of the results are discussed and summarized in §§ 6 and 7.

2. OBSERVATIONS

The $V$, $R$, and $I$ images used in this study were obtained at the prime focus of the Canada-France-Hawaii telescope using FOCAM and the LICK1 and LICK2 large format, 2k x 2k CCDs between 1991 April and 1993 March. The image scale for both LICK devices is $0.207$ per pixel so the full field of view of the CCD is $\sim 7'$ on a side. These images of high Galactic latitude “blank” fields were originally obtained for a survey of Population II halo stars (see Richer
Fig. 1.—$V + R + I$ image of the NF1 field with each bandpass normalized to an equivalent sky level. The areas of the CCD frame with cosmetic defects, saturated stars and very bright galaxies are masked out and not considered in the image analysis. Note the cosmetic differences between the LICK1 CCD (NF1) and the LICK2 device (NF2 and NF3; Figs. 2 and 3).

& Fahlman 1992) but are also useful for studying properties of the faint galaxy population. Three fields with north Galactic latitudes were observed and are dubbed NF1, NF2, and NF3. The right ascension and declination, and the corresponding Galactic coordinates, of the centers of these three blank fields are given in Table 1, along with the time of the observing runs. Fields were specifically chosen to have no observable objects on the Palomar Sky Survey photographs and a lack of any Zwicky clusters. Seeing for the frames used in the final summed images is uniformly excellent, ranging from 0:5 to 1". Good seeing is essential for acquiring deep images in a reasonable amount of exposure time. The filter bandpasses used, and the total exposure times and average seeing for the summed frames in each color and field, are summarized in Table 2. $V + R + I$ frames that demonstrate the total multiband exposure for each field are shown in Figures 1, 2, and 3.

3. DATA REDUCTION AND ANALYSIS

3.1. Image Preprocessing and Summation

Exposures of the fields were obtained typically for 900–1200 s, in between which the telescope would be dithered by $\approx 10^\circ$ in the cardinal directions such that program frames could be used to generate a sky flat in each bandpass, for flat-fielding purposes. Various programs from the IRAF\(^2\) package were used to do the preprocessing of the CCD

\(^2\) Image Reduction and Analysis Facility, a software system distributed by the National Optical Astronomy Observatories (NOAO).
images. All frames had the instrumental dc level, as determined from the CCD bias region, removed with the "linebias" routine and a median bias frame was then subtracted from each program frame to remove the pixel-to-pixel pattern. The pixel response function for each bandpass was determined by constructing a sky flat from the median of the corresponding dithered program frames. This "superflat" was generated using anywhere from 7 to 10 program frames, with this set including the data frames used for the final combined frame. Any artifacts caused by the use of a data frame in the making of its own flat field were dealt with by masking brighter objects, where these effects were most significant.

After the flat-fielding step, the individual frames were registered with the IRAF routine "imalign." The alignment interpolation was performed to the nearest pixel. There is a slight rotation (~1°) between the images obtained of NF2 in 1992 and 1993 and the IRAF routines "geomap" and "geotran" were used to align these data. Finally, an exposure weighted average of the aligned program frames in each field and for each bandpass was obtained with the IRAF routine "imcombine." CCD frames with poor seeing were not included in the average to avoid resolution degradation. The final combined frames were always found to be flat to within ≤1%. Image flatness is an important feature of our processed images in that it allows accurate faint galaxy photometry to be determined. Cosmic rays were also removed, as a final step, using the IRAF routine "cosmicrays." No obvious differences in the final images were found when the cosmic rays were removed before or after the combination of the individual data frames. Careful selection of the detection threshold and other parameters for the "cosmicrays" routine resulted in the removal of most of the cosmic rays from the final combined frames.

3.2. Calibrations

Standard star fields were observed on every night of each observing run. The fields used are summarized in Table 3,
Fig. 3.—$V + R + I$ image of the NF3 field with each bandpass normalized to an equivalent sky level, as in Fig. 1

| Observing Run  | Standards                        |
|----------------|----------------------------------|
| 1991 April     | M67, NGC 4147                   |
| 1992 June      | NGC 4147, SA 110, SA 113, M92    |
| 1993 March     | SA 98, M92, G12 – 43, RU 149     |

* Montgomery et al. 1993, Schild 1983.
* Christian et al. 1985.
* Landolt 1992.
* L. Davis, private communication.

The zero points in the calibration equations were very stable, not varying more than ~0.2 mag between observing runs. Color terms found in the calibration solutions for the three runs were accounted for by small offsets ($\leq 0.05$) based on a mean color for the faint galaxies or were disregarded due to the calculated coefficient being negligible. Stetson’s (1987, 1990) “DAOGROW” and “DAOPHOT” programs were used to determine accurate aperture corrections for the standard star frames. The only departure from standard calibration with the pertinent references.
techniques necessary was in the case of the NF2 $V$ and $I$
final frames, which were comprised of images from two dif-
f erent observing runs. A set of $\sim 20$ stars in the NF2 field
were chosen to be secondary standards. Magnitude offsets
between these secondary standards photometered on a
single frame from the 1992 June run, and the final, averaged
frames were determined. These small offsets ($\lesssim 0.05$ mag)
were found to be color independent and therefore could be
applied directly to the magnitudes obtained from the final
averaged $V$ and $I$ frames for NF2, to correct for the varia-
tions between the two observing runs. For more details of
the calibrations the reader is referred to Woods (1996).

3.3. Generating FOCAS Catalogs

The objects in our final CCD frames were detected and
analyzed using FOCAS (Jarvis & Tyson 1981) routines with
slight modifications to the standard analysis procedures,
some of which are outlined in Valdes (1983, 1993). The two
key input parameters for FOCAS are the detection thresh-
hold, given as a multiple of the sky variance ($\sigma_{\text{sky}}$), and
the minimum area for the objects detected, in numbers of pix-
els. After experimenting with these parameters and ensuring
that spurious detections were minimized, we adopted a
threshold of $2.5\sigma_{\text{sky}}$ and a minimum area which correspond-
ed to the seeing disk for the poorest resolved frame from
the three bandpasses taken of a given field (also see Steidel &
Hamilton 1993). Since we use conservative magnitude limits
in this study (see §3.4) where galaxy incompleteness is negli-
gible and the success of the galaxy detection is checked
thoroughly by eye, we are confident that the threshold and
minimum area parameters chosen are appropriate.

Two approaches were used for the initial detection of the
objects in our fields. The first technique was to use a
“master” frame ($V + R + I$), where the average frames
from each bandpass were normalized to a common sky level
in counts (Smail et al. 1994). An object list was generated
from the master frame detections, then the galaxy magni-
tudes were measured off of the average images in each
respective filter. This works quite well for fields where data
have been obtained with comparable magnitude limits in
the three filters (NF1). Note that $V \sim R + 0.5$ and
$R \sim I + 0.75$ (see galaxy color histograms in Fig. 7 below)
for galaxies with late-type morphologies within the redshift
regime that approximately corresponds to our magnitude
limits (Frei & Gunn 1994). Exceptions to this are ellipticals,
which become harder to detect since the 4000 Å break has
been redshifted beyond the $V$ filter at $z \gtrsim 0.5$. In fields
where the faint limits were not roughly equivalent from
filter to filter (NF2 and NF3), the second approach was to
have the initial detection of the objects done in each individ-
ual bandpass. The objects found in each filter were then
matched in master catalogs in order to provide color infor-
mation. In particular, the $R$-band data taken for NF2 and
NF3 were found to be deeper than the $V$ and $I$ data.

FOCAS programs are applied to either the master frame
of a field or the averaged frame in each filter (i.e., the
“detection” frame) to determine an initial catalog of
objects, with a slight modification. The detection algorithm
for FOCAS will find different numbers of objects, varying
by a few percent, depending on the orientation of the frame.
This variation in numbers is due to the line-by-line nature
of the detection algorithm and the fact that the threshold
for a particular line depends on the sky history from
the previous lines. We work around this problem by rotating
the detection frame through 90 degree increments and
matching the resultant four catalogs to produce a final
catalog. An object is included in the final catalog if it is
detected in all the catalogs from the four orientations. We
used the “builtin” FOCAS filter (1 2 3 2 1) for convolution
with the image under consideration during the detection
process to reduce the number of spurious objects in the final
catalogs. The final catalog is also filtered to remove detec-
tions of objects that lie within “masked” areas of the frame.
Masked regions include saturated stars, very bright gal-
axies, bad columns, vignetted corners, and other artifacts
that generate spurious detections. The masked area of
each field is typically only a few percent of the total number
of pixels in the final frame.

Following the application of the detection algorithm, the
sky values are determined for each frame using the standard
“sky” and “skycorrect” routines in FOCAS. No significant
development of the measured sky values on the orientation
of the frame was found. Detections listed in the final cata-
logs were split into individual objects and the magnitudes
were evaluated using the default FOCAS programs. Split-
ing of multicomponent groups into individual galaxies
(and stars) was easily accomplished for the separations over
which the angular correlation function is calculated, mostly
due to the uniformly excellent seeing of the data set. All
objects were split with confidence down to separations of
$\sim 1^\prime$, as in Woods et al. (1995). Measurement of galaxy
magnitudes is discussed further in the following section. A
point spread function (PSF) was determined from $\sim 15$–$20$
bright stars found in each field for each bandpass. The PSF
is used in the FOCAS object classification algorithm
resolution, to separate galaxies and stars from spurious
objects in the final master catalog. Separation of stars from
galaxies in the final samples was not done with the FOCAS
classification routine but with an approach outlined in §4.3.
Objects listed in the final catalogs were all checked by eye to
confirm their detection. A few spurious objects remained at
this juncture and were removed from further analysis, but
their numbers were small relative to the final galaxy sample.

3.4. Final Photometry

Total exposures for the three bandpasses, in a given field,
were obtained so as to be comparable in depth to maximize
the color information for the greatest number of objects. In
practice this is difficult to do at the telescope due to varying
seeing, along with weather and observing time constraints.
However, we obtained fairly uniform sampling of the
objects allowing us to determine $V$, $R$, and $I$ magnitudes for
$\sim 1000$ galaxies in each field, with the exceptions of the $V$
and $I$ data for NF2. Slightly lower numbers of faint objects
for the NF2 data were detected due to the data sets being
collected during two observing runs, which created a small
loss in area from slightly mismatched fields. Only the field
area that is coincident on all the data frames is included in
the final detection frames, so that uniform magnitude-
limited samples are generated.

Magnitudes were evaluated for each bandpass and field
yielding final lists of isophotal and aperture magnitudes,
and colors for all the objects. Aperture magnitudes were
found to be a more reliable measure of the total brightness
of the faint galaxies than isophotal magnitudes. Also, the
isophotal magnitude limit was found to be unrealistically
faint. However, the use of a small fixed aperture is not
appropriate for the galaxies in the sample that have a sig-
significant angular extent. Hence, “hybrid” magnitudes are adopted: an aperture of 3'' in diameter was used for the galaxies with an average diameter \( \leq 3'' \) and isophotal magnitudes were used for galaxies with characteristic sizes larger than this. Isophotal and aperture magnitudes are in good agreement at the magnitude range where the transition between the two measures occurs, typically having differences \( \leq 0.1 \) mag.

Isophotal magnitudes are usually used for galaxies that are up to 3 mag below the bright magnitude limit \( (V \sim 20-23, R \sim 20-23 \text{ and } I \sim 19-22) \) where galaxies are more than 3'' in mean diameter. Since the majority of the galaxies in any of our samples is within 2 mag of the faint magnitude limit \( (V > 23, R > 23, \text{ and } I > 22) \) the galaxies measured with isophotal magnitudes are a small contribution to the final sample size. The isophotal magnitudes used are not the “total” magnitudes which FOCAS generates (Valdes 1983) since we prefer to avoid doubling the isophotal area, from the initially determined isophote, for the flux measurements. Considering the small numbers of galaxy magnitudes measured with isophotes this choice does not have a significant effect on the final magnitude-limited galaxy samples. Apertures of 3'' were chosen since they were determined to be large enough to contain most of the flux from the majority of the faint galaxies. The adopted aperture size of 3'' corresponds to a physical scale of \( \sim 11 \, h^{-1} \) kpc and \( \sim 13 \, h^{-1} \) kpc for redshifts of \( z = 0.5 \) and 1.0, respectively, using \( h = H_0/100 \) km s\(^{-1}\) Mpc\(^{-1}\) and \( q_0 = 0.5 \). See \$ 2.3 of Lilly, Cowie, & Gardner (1991) for some further discussion of the benefits of aperture photometry for faint galaxies, as opposed to using isophotal photometry.

Number counts, in \( V, R, \) and \( I \), are given for the three fields in Figures 4, 5, and 6, respectively. Bright stars have been removed from these counts in the manner outlined in \( \S \) 4.3 and hybrid magnitudes are used to calculate these counts. The slopes of the galaxy counts are listed in Table 4. There is good agreement between the slopes determined in different fields for a given bandpass. These are encouraging results since we want to calculate \( \omega(\theta) \) by averaging the galaxy clustering behavior in NF1–3, thus requiring field-to-field uniformity. To avoid crowding the plots in Figures 4–6 with a plethora of data points from other studies we...
have compared our $R$ number counts normalization and slope to the compilation of Metcalfe et al. (1991) and the counts of Hudon & Lilly (1996, hereafter HL), both plotted in the latter study’s Figure 1. We find excellent agreement with the $R$ counts of HL and Metcalfe et al. Checking our $V$ and $I$ data is more problematic due to a scarcity of published number counts in these bandpasses. Our $V$ counts are in good agreement with the observations shown in Figure 2 of Smail et al. (1995) except we observe no break at $V \sim 24.25$, but this is not surprising considering our magnitude limit is $V \sim 25$. The $I$ counts from the current study have a slightly steeper slope and different normalization than the Smail et al. data but are consistent with the Lilly et al. (1991) number counts. The source of the difference with Smail et al. is not clear, however, they suggest that the discrepancy could be due to other workers choosing fields that are devoid of bright galaxies therefore biasing the count slope to be steeper. Nevertheless, we find our number counts to be in generally good agreement with other groups’ observations.

From the number counts it is easily seen that conservative magnitude limits for the data are $V \sim 25$, $R \sim 25$, and $I \sim 24$, except the NF1 $R$ data, which have a limit of $\sim 24.5$. This latter limit is more comparable to the $V$ and $I$ limits for galaxies at intermediate redshifts ($z \sim 0.5—0.7$). To test the completeness levels of galaxy detection at these magnitude limits we checked the number of “dark” galaxies detected by FOCAS (or negative fluctuations on the CCD) in each field and bandpass. The excess of above-sky objects to the number of below-sky objects detected can be used as an unbiased estimate of the number of faint objects at that magnitude (Valdes 1983). All of our samples were found to be essentially $\sim 100\%$ complete at the magnitude limits that are listed above, confirming the conservative nature of our faintness limits. Photometric errors were calculated to be typically $\lesssim 0.1$ mag. for galaxies with $V < 24$, $R < 24$, and $I < 24$, increasing to as much as $\sim 0.3$ mag for fainter galaxies. In Table 5, the number of objects found within the given magnitude limits are tabulated, along with the effective field areas. All of these objects comprise the magnitude-limited samples that are used for angular correlation analysis in § 5.

Color-selected samples were obtained by matching the catalogs generated for a given field for the two bandpasses in the required color. The magnitude limit for each color-selected sample was chosen to be the limit of the bandpass (either $V \sim 25$, $R \sim 25$, or $I \sim 24$, except NF1 where it is $R \sim 24.5$) with the largest number of galaxies, thereby maximizing the accuracy of the clustering measurements. This approach introduces some galaxies into the sample that are beyond the magnitude limit for the bandpass with the smaller sample size, in order to maintain a high level of matching. Colors are calculated using “hybrid” magnitudes, but the number of galaxies measured with isophotal magnitudes is small enough such that there is little effect on our final results. This was checked by calculating the results using galaxy samples that contained only aperture magnitudes. No attempts were made to convolve the two bandpass images to the poorest seeing before the colors were determined since the majority of galaxies easily fit into the 3’ apertures adopted and the seeing values in different bandpasses for the same field are comparable (Table 2).

High percentages (typically $\geq 90\%$) of the galaxies matched between filters, particularly for the NF1 field where the magnitude limits in $V$, $R$, and $I$ were of similar depth. A 5%—10% decrease in the match success rate was observed for the faintest magnitude bins. These faint galaxies are not detected in all three filters partly because of color effects and because the counts are beginning to become incomplete due to fluctuations in the background. Histograms of the colors of the galaxies found in the three fields are shown in Figure 7. $(V - R)$, $(R - I)$, and $(V - I)$ galaxy colors are shown in the plot panels from top to bottom for NF1 (solid line), NF2

| Field | $V$ | $R$ | $I$ |
|-------|-----|-----|-----|
| NF1   | 0.41 ± 0.01 | 0.36 ± 0.01 | 0.32 ± 0.01 |
| NF2   | 0.42 ± 0.02 | 0.35 ± 0.01 | 0.34 ± 0.02 |
| NF3   | 0.46 ± 0.02 | 0.39 ± 0.02 | 0.33 ± 0.02 |

### Table 5

| Filter     | Number of Objects (Magnitude Range) |
|------------|-------------------------------------|
| NF1 (Effective Area: $\sim 0.01064$ deg$^2$)… |
| $V$        | 355 (20–24) | 590 (20–24.5) | 935 (20–25) |
| $R$        | 574 (20–24) | 878 (20–24.5) |… |
| $I$        | 486 (19–23) | 706 (19–23.5) | 996 (19–24) |
| NF2 (Effective Area: $\sim 0.01026$ deg$^2$)… |
| $V$        | 253 (20–24) | 425 (20–24.5) | 699 (20–25) |
| $R$        | 407 (20–24) | 610 (20–24.5) | 913 (20–25) |
| $I$        | 319 (19–23) | 496 (19–23.5) | 709 (19–24) |
| NF3 (Effective Area: $\sim 0.01111$ deg$^2$)… |
| $V$        | 312 (20–24) | 543 (20–24.5) | 908 (20–25) |
| $R$        | 588 (20–24) | 930 (20–24.5) | 1482 (20–25) |
| $I$        | 439 (19–23) | 658 (19–23.5) | 992 (19–24) |

![Fig. 7. Histograms of galaxy colors for NF1, NF2, and NF3 with the number of galaxies observed as a function of $(V - R)$, $(R - I)$, and $(V - I)$ plotted in the top, middle, and bottom panels, respectively. The NF1, NF2, and NF3 fields are indicated by solid-line, dashed-line, and dotted-line histograms.](image-url)
4. MEASURING THE ANGULAR CORRELATION FUNCTION

4.1. Estimator

If one considers two differential elements of solid angle on the sky, \( d\Omega_1 \) and \( d\Omega_2 \), then the joint probability, \( dP \), that galaxies will occupy the two elements with an angular separation of \( \theta \) can be written as

\[
dP = n^2 [1 + \omega(\theta)] d\Omega_1 d\Omega_2 ,
\]

where \( \omega(\theta) \) is the angular correlation function and \( n \) is the mean surface density of galaxies. A random (Poisson) distribution of galaxies yields \( \omega(\theta) = 0 \) for all \( \theta \). Therefore, \( \omega(\theta) \) is simply a measure of the number of galaxy pairs observed at a given separation projected on the sky, normalized by the number of galaxy pairs expected if the galaxies are randomly distributed. The traditional estimator for \( \omega(\theta) \) (Peebles 1980), where equal numbers of galaxies and random points are considered, is of the form

\[
\hat{\omega}(\theta) = \frac{DD(\theta)}{RR(\theta)} - 1 ,
\]

where \( DD(\theta) \) and \( RR(\theta) \) represent the number of data-data and random-random pairs at the angular separation \( \theta \) (hereafter pair symbols will implicitly be assumed to be functions of \( \theta \) and the hat symbol is used to denote an estimate of a function). Another estimator includes a cross-correlation of the data and random objects:

\[
\hat{\omega}(\theta) = \frac{2DD}{DR} \left( \frac{N_K}{N_R - 1} \right) - 1 ,
\]

(see EBKTG, Infante & Pritchet 1995, and references therein), where \( N_R \) and \( N_K \) are the number of galaxies and random objects, respectively. The estimators in equations (2) and (3) have greater than Poissonian variance, so, to minimize the noise we adopt the estimator suggested by Landy & Szalay (1993, hereafter LS; also see Hamilton 1993):

\[
\hat{\omega}(\theta) = \frac{(DD - 2DR + RR)}{RR} .
\]

For a given galaxy sample, 100 files of random positions, each containing the same number of "galaxies" as those observed, were generated, yielding \( \sim 0.5 - 1.0 \times 10^6 \) random objects. An increase in the number of random position files from 100 to 1000 did not significantly improve the accuracy of the final \( \omega(\theta) \) estimation but substantially increased the computing time. The random position files were created using the same detection mask that was used for the galaxy image (in all three bandpasses for a given field) to block out saturated objects and areas of the CCD with defects and vignetting. A mean of the 100 random files cross correlated with the particular galaxy sample yielded DR in the estimator. In addition, by averaging the files, one can calculate the probabilities of obtaining a pair and triplet at a separation \( \theta \) [LS refers to these probabilities as \( G_1(\theta) \) and \( G_2(\theta) \), respectively], which are quantities required to calculate \( \omega(\theta) \) and the errors associated with the estimator.

To check the clustering estimator in equation (4), we also determined counts in cells using

\[
\hat{\omega}(\theta) = \frac{\langle N_i N_j \rangle}{\langle N_i \rangle \langle N_j \rangle} - 1 ,
\]

(5)

with the number of galaxies in cells \( i \) and \( j \) denoted by \( N_i \) and \( N_j \), and the angular brackets representing an average of all the cells with an angular separation within the bin \( \theta \pm \delta \theta \). Images were divided into square cells, 5" on a side. Excellent agreement was obtained between the data pairs (eq. [4]) and cell counts estimators, so, in the remainder we only refer to the data pairs approach.

4.2. The Clustering Model

The standard model for the angular correlation function is

\[
\omega(\theta) = A_\theta \theta^{-\delta} ,
\]

(6)

where \( \delta \) has been found to range from \( \sim 0.6 \) to 0.8 for faint samples. The value of \( \delta \) could be dependent on the angular scales \( (\theta) \) probed or the magnitude limit of the galaxy sample (Maddox et al. 1990, BTBJ, and Neuschaefer & Windhorst 1995). If we want to use the estimated \( \omega(\theta) \) projected on the sky to determine the spatial two-point correlation function, we must adopt a model that includes possible clustering evolution with redshift \( (z) \). The conventional model for the spatial correlation function is

\[
\xi(r, z) = \left( \frac{r}{r_0} \right)^{\gamma} (1 + z)^{-(3 + \alpha)} ,
\]

(7)

where evolution with redshift is parameterized by \( \epsilon \), \( r_0 \) is the proper length, and \( r_0 \) is the correlation length at \( z = 0 \) (see Phillipps et al. 1978, Peebles 1980, EBKTG, Infante & Pritchet 1995 and Hudon 1995). If the redshift distribution \( (dN/dz) \) is known for the magnitude limits for which the angular correlation function has been measured, the relationship between \( \omega(\theta) \) and \( \xi(r, z) \) is determined by an integral known as Limber's equation (see, e.g., Peebles 1980 and HL):

\[
\omega(\theta) = C r_0^\gamma \int_0^\infty \frac{dN}{dz} \left[ \frac{dN}{dz} \right]^{-\gamma - 1} (1 + z)^{-3 - \alpha} dz ,
\]

(8)

(\( D(z) \) is the angular diameter distance defined as

\[
D(z) = \frac{c}{H_0} q_0 z + (q_0 - 1) \sqrt{1 + 2q_0 z - 1} ,
\]

(9)

with \( g(z) \) and \( C \) given by

\[
g(z) = \frac{c}{H_0} \left[ (1 + z)^{2(1 + 2q_0 z)^{1/2}} \right]^{-1} ,
\]

(10)

and

\[
C = \sqrt{\pi^{-1}} \frac{\Gamma((\gamma - 1)/2)}{\Gamma(\gamma/2)} ,
\]

(11)
with \( q_0, H_0, \) and \( \Gamma \) being the deceleration parameter, Hubble constant and Gamma function, respectively. Note that equation (7) and Limber’s equation (eq. [8]) gives \( \delta = \gamma - 1 \). If \( \gamma \approx 1.8 \), the value derived from surveys of bright, nearby galaxies (Davis & Peebles 1983), in the case of clustering which is fixed in comoving coordinates then \( \epsilon = \gamma - 3 = -1.2 \). “Stable clustering,” where the clustering is fixed in proper coordinates, is the result when \( \epsilon = 0 \). If \( \epsilon > 0 \), then there is a growth in the clustering with redshift in proper coordinates. Using an extrapolated redshift distribution \((dN/dz)\) from the CFRS (Lilly et al. 1995a), with the same faint magnitude limit as our galaxy sample \((I \leq 24)\), we can solve Limber’s equation for assumed values of \( q_0, H_0, \gamma, \) and \( \epsilon \). The growth of clustering \( (\epsilon) \) can then be estimated by comparing the models with the observations of \( \omega(\theta) \) (see § 5.3).

4.3. Star Removal

The number counts of objects at faint magnitudes are dominated by galaxies but at bright limits the stellar component makes a significant contribution. To correct for this stellar contamination we plot a shape parameter versus the aperture magnitudes for all the detected objects. This shape parameter is simply the difference between the “core” and aperture magnitude of an object, where the former is the magnitude corresponding to the flux incident on the inner \(3 \times 3\) pixels. The core and aperture magnitude difference is a measure of the object’s light concentration and is analogous to Kron’s (1980) \( r^-2 \) statistic, which is proportional to the half-light radius.

A plot of the core-aperture versus aperture \((3')\) magnitudes for \( V, R, \) and \( I \) in NF2 and \( I \) in NF3, is given in Figure 8. Star-galaxy separation diagrams for NF2 and NF3 with the difference between the “core” and aperture magnitudes plotted on the ordinate and just the latter on the abscissa. The field and bandpass are shown in the upper left corner of each panel. Magnitude limits are marked with dashed lines, and the stellar sequences are demarcated by solid line rectangular boxes. Note the “plume” of galaxies seen in the NF3-I panel, at the upper right of the data near the magnitude limit, is caused by spurious noisy objects yet to be removed from the galaxy catalog.

Fig. 8.—Star-galaxy separation diagrams for NF2 and NF3 with the difference between the “core” and aperture magnitudes plotted on the ordinate and just the latter on the abscissa. The field and bandpass are shown in the upper left corner of each panel. Magnitude limits are marked with dashed lines, and the stellar sequences are demarcated by solid line rectangular boxes. Note the “plume” of galaxies seen in the NF3-I panel, at the upper right of the data near the magnitude limit, is caused by spurious noisy objects yet to be removed from the galaxy catalog.
A correction must be made for the integral constraint, which is due to the estimation of the density of galaxies, at a given magnitude limit, with a bounded, finite sample (Peekes 1980). This bias has the effect of reducing the amplitude of $\omega(\theta)$. Following BSM and LS we calculate the integral constraint ($\omega_{i1}$) using

$$\omega_{i1} = \frac{1}{\Omega} \int \int \omega(\theta) d\Omega_1 d\Omega_2 , \tag{12}$$

with $\Omega$ representing the solid angle of the masked field. We also assume the angular correlation function has the functional form given in equation (6). Integral constraints are calculated for each field and for power laws with $\delta$ ranging from 0.5 to 0.9, resulting in $\omega_{i1} \sim 0.084 A_0 - 0.01 A_\omega$. For $\delta = 0.8$, the integral constraints are determined to be $\omega_{i1} \approx 0.0195 A_\omega$, $0.0199 A_\omega$, and $0.0193 A_\omega$ for NF1, NF2, and NF3, respectively. The values are comparable since the field sizes and geometries are similar. This correction is significant compared to the small amplitudes of $\omega(\theta)$ measured from the faint galaxy samples (see §5).

4.5. Star Dilution Correction

Since stars have not been removed from the photometric samples fainter than $V$, $R$, or $I \sim 22$, a correction must be made to the amplitude of $\omega(\theta)$ to account for the stellar component present down to the magnitude limit. The number of stars expected at faint magnitude limits is taken from the model of Bahcall & Soneira (1980). The three fields in this study were chosen to be at high Galactic latitudes where stars are relatively scarce and therefore the stellar dilution corrections are small. The amplitude of $\omega(\theta)$ after the correction for stellar dilution is made is given by

$$A_{\omega}^{se} = \left( \frac{N_{obj}}{N_{obj} - N_s} \right)^2 A_\omega , \tag{13}$$

where $N_{obj}$ is the number of objects used to calculate $\omega(\theta)$, $N_s$ is the number of stars predicted by the Bahcall & Soneira model and $A_\omega$ is the “raw” amplitude of $\omega(\theta)$ before any corrections or weighting have been applied. Stellar dilution correction terms calculated for the various magnitude-limited samples are listed in Table 7 below.

4.6. Combining Fields, Error Analysis, and Fitting

Since the use of just one of the three observed deep fields leads to a determination of $\omega(\theta)$ that is of low accuracy, a strategy must be adopted to combine the data sets and to calculate the average or “final” $\omega(\theta)$ and the appropriate errors. We follow a similar approach to that of Neuschaefer et al. (1995, hereafter NRGCI), where the average $\omega(\theta)$ for a given bin ($\theta$) is calculated using

$$\hat{\omega}_{\text{fin}}(\theta) = \frac{\sum \eta_i \hat{\omega}^{IC}_{i}(\theta)}{\sum \eta_i} , \tag{14}$$

where the summations extend over fields $i = 1, 2, 3$, and where the $\eta_i$ are the weights for each field and include the stellar dilution correction and the galaxy number densities. We use the number densities of objects as weights since the total areas of the three fields are slightly different. The “IC” superscript on the $\omega(\theta)$ estimates for each field denotes that corrections have been made for the integral constraint. Corrections for higher order correlations (e.g., three-point correlation function) are disregarded since they are negligible relative to the values of the $\hat{\omega}^{IC}_{i}(\theta)$ calculated in the individual fields.

Although the estimator for $\omega(\theta)$ that we use has been shown to give Poissonian variance for uncorrelated data by LS, it does not necessarily follow that it behaves this way for correlated data. This was first pointed out by Bernstein (1994), who also emphasized that most authors do not properly account for the interdependence between the various bins when estimating the uncertainty in their results. We calculate errors using a scheme outlined by Fisher et al. (1994) where the covariance matrix for a particular estimate of $\omega(\theta)$ is determined with bootstrap resampling (Barrow, Bhavsar, & Sonoda 1984). Alternatively, Bernstein (1994) derives an approximate analytical expression for the covariance matrix, but the model fitting procedures in either study are essentially equivalent. As in NRGCI, for a given field and magnitude-limited sample, resampled estimates of $\hat{\omega}^{IC}_{i}(\theta)$ are calculated by applying the estimator of equation (4) to a resampled list of galaxies with the same number of objects as the original, real sample. The resampled list is generated by randomly selecting galaxies from the original list with replacement, such that a galaxy can be chosen anywhere from zero up to several times. For each magnitude-limited sample, 50 bootstrap-resampled estimates of $\hat{\omega}_{\text{fin}}(\theta)$ are made by averaging resampled estimates of $\hat{\omega}^{IC}_{i}(\theta)$ calculated for the three fields. The final bootstrap errors for the different angular separation bins are simply given by the variance of the resampled estimates of $\hat{\omega}_{\text{fin}}(\theta)$. Finally, a covariance matrix is generated so that the power-law model (eq. [6]) can be properly fit to the clustering observations for each magnitude-limited sample, following the technique described by Fisher et al. (1994, Appendix A).

Since the galaxy samples are fairly small (see Table 5), we choose to fix the power-law exponent ($\delta$) and only let the amplitude ($A_\omega$) vary when fitting the model (eq. [6]) to the
The \( \chi^2 \) minimization is analytic with one linear parameter in the model. The power-law exponent, \( \delta \), has been measured to range from \( \sim 0.6 \) to \( \sim 0.9 \) (Neuschaefer & Windhorst 1995, and references therein) for galaxies at faint magnitude limits. Accordingly, each \( \hat{\delta}(\theta) \) calculated for a magnitude-limited sample is fit with power laws between \( \delta = 0.5 \) to \( \delta = 0.9 \), in 0.1 increments. The \( \chi^2 \) statistic calculated for each fit gives an idea of what the most appropriate value for \( \delta \) is, although the measurement is not well constrained. Our data favor larger values of \( \delta (\delta \sim 0.8--0.9) \), so we fix \( \delta = 0.8 \) to ease comparison with other studies. Table 6 illustrates the relative insensitivity of the final amplitudes to fits with power laws having different values for \( \delta \), in this case for the faintest magnitude-limited samples in \( V \), \( R \), and \( I \). The \( \chi^2 \)-values, which are for five degrees of freedom, decrease as the power law approaches \( \delta = 0.8--0.9 \) but for \( \delta > 1.0 \) the errors in the amplitude fit increase dramatically. For decreasing values of \( \delta \) the integral constraint increases (§ 4.4) and this is the primary reason the \( \chi^2 \) statistic increases dramatically, as listed in Table 6. Given the limited statistics of our sample, fixing \( \delta = 0.8 \) seems to be the best approach but it should be emphasized that the errors for the fitted amplitudes using this technique are probably underestimates. Further details of the results of the model fitting are discussed in § 5.1 below.

For each magnitude-limited sample in a particular field, \( \hat{\delta}(\theta) \) is calculated for angular separations ranging from \( 10^\circ--126^\circ \), within six equally spaced logarithmic bins. The binning and angular separation range were carefully chosen to optimize the measurement of \( \omega(\theta) \) given the available imaging data. The upper limit for \( \theta \) was chosen to be roughly one-third of the angular extent of the smallest field, thereby avoiding border effects. A lower limit of 10\( ^\circ \) yielded error bars for the smallest angular separation bin that were roughly comparable to those obtained for bins with a larger \( \theta \). A discussion of close galaxy pairs with smaller angular separations (\( \leq 10^\circ \)) is given in Woods et al. (1995).

5. ANGULAR CORRELATION FUNCTION RESULTS

5.1. Magnitude-limited Samples

Measurements of the angular correlation function for the magnitude-limited samples defined in § 3.4, for \( V \), \( R \), and \( I \), are presented in Figures 9, 10, and 11 respectively. The correlation amplitudes generally decrease over the small range of magnitudes probed, most obviously with the \( R \) data. The solid lines in Figures 9–11 are fits of the model, \( \omega(\theta) = A_0 \theta^{-0.8} \), to the data and the errors are calculated using bootstrap resampling, as described in the previous section. Amplitudes, \( A^{\text{fit}}_m \), measured from the fits to \( \hat{\omega}(\theta) \) for the various magnitude ranges are listed in Table 7 and are scaled for angular separations (\( \theta \)) given in arcseconds. The stellar dilution corrections used for the galaxy samples are probably underestimates. Further details of the results of the model fitting are discussed in § 5.1 below.

![Figure 9](image1.png)

**Fig. 9.**—Measurements of the angular correlation function using \( V \)-band data from NF1, NF2, and NF3, for the listed magnitude ranges. Error bars are calculated using bootstrap resampling. The solid lines represent the standard model fit to the data assuming \( \delta = 0.8 \). See text for details.

![Figure 10](image2.png)

**Fig. 10.**—Same as in Fig. 9, but for the \( R \) data.

| \( \delta \) | \( 20 \leq V \leq 25 \) | \( 20 \leq R \leq 25 \) | \( 19 \leq I \leq 24 \) |
|---|---|---|---|
| \( 0.5 \) | \( 0.653 \pm 0.024 \) | \( 0.686 \pm 0.022 \) | \( 0.689 \pm 0.020 \) |
| \( 0.6 \) | \( 0.596 \pm 0.036 \) | \( 0.640 \pm 0.033 \) | \( 0.641 \pm 0.031 \) |
| \( 0.7 \) | \( 0.546 \pm 0.052 \) | \( 0.605 \pm 0.049 \) | \( 0.603 \pm 0.045 \) |
| \( 0.8 \) | \( 0.508 \pm 0.077 \) | \( 0.583 \pm 0.072 \) | \( 0.578 \pm 0.067 \) |
| \( 0.9 \) | \( 0.485 \pm 0.111 \) | \( 0.581 \pm 0.104 \) | \( 0.570 \pm 0.097 \) |

**Table 6**

| \( \delta \) | \( 20 \leq V \leq 25 \) | \( 20 \leq R \leq 25 \) | \( 19 \leq I \leq 24 \) |
|---|---|---|---|
| | \( A^{\text{fit}}_m \) | \( \chi^2 \) | \( A^{\text{fit}}_m \) | \( \chi^2 \) | \( A^{\text{fit}}_m \) | \( \chi^2 \) |
| \( 0.5 \) | \( 0.653 \pm 0.024 \) | 49.3 | \( 0.686 \pm 0.022 \) | 62.6 | \( 0.689 \pm 0.020 \) | 67.5 |
| \( 0.6 \) | \( 0.596 \pm 0.036 \) | 26.9 | \( 0.640 \pm 0.033 \) | 36.7 | \( 0.641 \pm 0.031 \) | 37.9 |
| \( 0.7 \) | \( 0.546 \pm 0.052 \) | 13.7 | \( 0.605 \pm 0.049 \) | 20.9 | \( 0.603 \pm 0.045 \) | 20.2 |
| \( 0.8 \) | \( 0.508 \pm 0.077 \) | 6.6 | \( 0.583 \pm 0.072 \) | 11.9 | \( 0.578 \pm 0.067 \) | 10.5 |
| \( 0.9 \) | \( 0.485 \pm 0.111 \) | 3.1 | \( 0.581 \pm 0.104 \) | 7.1 | \( 0.570 \pm 0.097 \) | 5.4 |
are also listed in Table 7. To demonstrate the scatter of the uncorrected estimates of \( \omega(\theta) \) between the three fields, the “raw” angular correlation functions are plotted for the \( I \)-band data in Figure 12. The circles, triangles, and squares are data points measured from the NF1, NF2, and NF3 fields, respectively. NF1 data points have been offset 0.01 dex to the left, while the NF3 data points have been shifted the same amount to the right, for the purpose of clarity. These measurements have not been corrected for the integral constraint or for stellar dilution (see Table 7). Better agreement is observed between the three fields at fainter magnitude limits, where the samples of galaxies are the largest. Note the ordinate scale covers a smaller range than in Figure 11.

![Figure 11](image1.png)

**Figure 11.**—Same as in Fig. 9, but for the \( I \) data.

![Figure 12](image2.png)

**Figure 12.**—Comparison of “raw” angular correlation functions measured for the listed magnitude ranges in each field, in the \( I \) bandpass. The circles, triangles, and squares denote the NF1, NF2, and NF3 data points, respectively. NF1 data are offset 0.01 dex to the left, while the NF3 data points have been shifted the same amount to the right, for the purpose of clarity. These data have not yet been corrected for either the integral constraint or for dilution due to stars. The improvement of the agreement of the \( \omega(\theta) \) estimates in the three fields toward fainter magnitude ranges is typical for the three bandpasses and is due to the larger numbers of galaxies in the faint samples. Note that the ordinate range of is smaller than Figure 12 to better illustrate the similarities between the three fields. In Figures 9–11 it can be seen that the faintest sample for each bandpass appears to have a slightly flatter distribution than the assumed \( h^{-0.8} \) power law. Since there is good agreement between the clustering measurements made in the three fields at faint limits this effect could be real and possibly connected with the flattening of the power-law slope of \( \omega(\theta) \) observed by Neuschaefer & Windhorst (1995). As discussed in the previous section, our galaxy samples are of insufficient size to accurately determine both the amplitude \( (A_w) \) and the power-law slope \( (\delta) \) simultaneously. Assuming \( \delta < 0.8 \) introduces a larger integral constraint correction for the data plotted in Figures 9–11 and thereby increases the \( \chi^2 \) value for the fit. The adoption of \( \delta = 0.8 \) is to ease the comparison with other studies but it is still typically the best fit when all the corrections to the data are made. However, it should be noted that there is some evidence for a decrease in the value of \( \delta \) at fainter limits and this should be investigated with larger samples of faint galaxy photometry.

The values of the angular correlation functions measured in \( V \), \( R \), and \( I \) at the separation of \( \theta = 1' \) are plotted in Figure 13. The \( R \)-band data gives the strongest clustering

**Table 7**

| Band | Magnitude Ranges | \( A_w \) | \( \chi^2 \) |
|------|------------------|----------|----------|
| V    | 20 ≤ \( V \) ≤ 24 | 0.550 ± 0.233 | 0.508 ± 0.077 |
|      | 20 ≤ \( V \) ≤ 24.5 | 0.527 ± 0.141 |
|      | 20 ≤ \( V \) ≤ 25   | 0.527 ± 0.141 |
| R    | 20 ≤ \( R \) ≤ 24   | 0.504 ± 0.176 | 0.578 ± 0.067 |
|      | 20 ≤ \( R \) ≤ 24.5 | 0.627 ± 0.116 |
|      | 20 ≤ \( R \) ≤ 25   | 0.627 ± 0.116 |
| I    | 19 ≤ \( I \) ≤ 23   | 1.083 ± 0.176 | 1.065 ± 0.067 |
|      | 19 ≤ \( I \) ≤ 23.5 | 1.075 ± 0.116 |
|      | 19 ≤ \( I \) ≤ 24   | 1.075 ± 0.116 |
signal but the variations are roughly comparable to the
errors. Lilly et al. (1995a) have measured the redshift dis-
btribution of faint galaxies selected to a magnitude limit of
$I \approx 22$ and also make extrapolations for $N(z)$ to limits as
faint as those obtained for the photometry in this work.
With these redshift distributions and the $I$-band angular
correlation analysis by Lidman & Peterson (1996, hereafter
LP) at brighter magnitude limits, we use our $\omega(\theta)$ measurements
in $I$ to constrain model parameters for the spatial
correlation function in $\S$ 5.3. Finally, since there have been
many recent correlation studies of faint galaxies in $R$
bandpasses, these provide a comparison for the clustering
detected with $R$-filter data in this study.

In Figure 14, the angular correlation function normalized
to $\theta = 1^\circ$, using the standard power-law model with $\delta = 0.8$,
is plotted as a function of the $R$ magnitude limit. The
measurements by different groups are denoted by the various
symbols, which are keyed to the authors’ initials and year of
the particular paper. The $\omega(\theta)$ amplitudes determined in this
work are given by the solid circles for $R = 20–24, 20–24.5,$
$20–25$. Magnitude transformations for BSM were made
using their assumption of $R \sim r - 0.55$. For Couch et al.
(1993) and others the conversions given by Yoshii, Peterson,
& Takahara (1993) and Roche et al. (1993) yield $R$
magnitudes from the original $VR$ and $r_p$ values. The only observation
plotted in Figure 14 that was not taken with a red filter is that of MSFR.
Since the MSFR result is the $\omega(\theta)$ amplitude with the faintest magnitude limit yet measured
from the ground, it is interesting to include it for comparison
using the approximate relationship $B_{\text{CCD}} \approx R + 1$.

It is notable that the $\omega(\theta)$ values presented in this work
form a smooth continuation of the previous observations
made by Infante & Pritchet (1995) and HL for $R = 21–23.5$,
where the latter study used the same $R$ filter as the current
observations. Our data agree reasonably well with the
EBKTG data point and very well with the overlapping
observations of Roche et al. (1996). The largest discrepancy
with this study is seen with BSM’s $\omega(\theta)$ measurements
where our clustering amplitudes are observed to be factors of $\sim 2–3$ larger. A possible explanation for part of this
difference is that the BSM field is at low Galactic latitude
($b \approx 35^\circ$) typically requiring larger stellar contamination
corrections than our three high Galactic latitude fields.
Another possibility is that our clustering amplitude errors
are underestimate since they are the fitting errors for a $\theta^{-0.8}$ power law. Nevertheless, the clustering measurements
in this study are in agreement with Roche et al. (1996), who
in turn agree with the BSM results, so there is a reasonable
level of consistency between studies in the $R \sim 24–25$
magnitude range. With three fields the current work is less sus-
ceptible to variations in general clustering behavior induced
by large scale structure.

The fact that our $VR$ and $I$-band estimates of $\omega(\theta)$ (Fig. 13)
do not show well defined decreases with magnitude, as the
$R$-filter data does, is not surprising due to the poorer sta-
tistics of the $V$ and $I$ filters for the two brightest magnitude bins (see Table 5). Therefore, the apparent flatness of the
clustering amplitudes with magnitude for the $V$ and $I$
galaxy samples should not be interpreted as a strong trend
but merely a clustering measure over a small range of mag-
nitudes (note the galaxy samples are cumulative toward
fainter magnitudes, not differential). Since clustering over a
significant magnitude range cannot be tracked with just the
data from this study, other studies must be included for a
proper analysis of galaxy clustering evolution. This is done for
the $I$-band data in $\S$ 5.3.

Our Figure 14 follows the format of Figure 2 presented in
BSM. Even with the additional results included from recent
studies, there still is not general agreement on the precise
slope of the monotonic decrease in clustering amplitude
with limiting magnitude. No clear indication that the ampli-

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**Fig. 13.**—Amplitudes of the angular correlation function calculated for
a separation of $1^\prime$ for $V$, $R$, and $I$, assuming a power law with $\delta = 0.8$.
Measurements for the $V$-, $R$-, and $I$-band data are shown in the left, center,
and right panels, respectively.

**Fig. 14.**—Amplitudes of the angular correlation function normalized to
1 degree assuming $\delta = 0.8$ for this work and other studies from the liter-
ature, in the $R$ band. Details of the magnitude transformations used for the
different observations are given in the text. Each symbol is listed with the
initials of the authors’ names and the year of the study it denotes.
tude is starting to level off at faint magnitudes is observed, as would be expected in some merger models of galaxy evolution (Carlberg & Charlot 1992) or if there was a magnification bias from weak gravitational lensing (Villumsen et al. 1993). Roche et al. (1993) and Roukema & Yoshii (1993, hereafter RY) have also found merging-model clustering behavior to be inconsistent with current measurements of \( \omega(\theta) \). MSFR claim that the amplitude of \( \omega(\theta) \) flattens for \( B \) data at about the same magnitude where the slope of the number counts flattens \( (B \sim 25 \text{ or } R \sim 24) \), which they attribute to an effective redshift cutoff for the galaxies. The data from various groups plotted in Figure 14 shows that any flattening in the clustering amplitude with magnitude is not that well constrained as yet, especially considering the inherently large random and systematic errors that plague the measurement of \( \omega(\theta) \) at faint limits.

One cannot rule out the viability of all merger models of galaxy evolution with current observations but models with extreme amounts of merging or “maximal merging” as in RY can be shown to be inconsistent with measurements of \( \omega(\theta) \). (Note that merging is considered “maximal” with the assumption that every galaxy merges when its dark halo merges.) See RY for more details on the nature of their galaxy merging models. The clustering measurements for the increasingly faint magnitude-limited samples in the \( V \) bandpass that are listed in Table 7 and shown in Figure 13 (with respective median magnitudes of \( V_{\text{median}} = 23.33, 23.84, 24.30 \)) have corresponding predictions from the RY merger models which are log \( [\omega(\theta = 1', \delta = 0.8)] \simeq -1.06, -1.14, \text{ and } -1.25 \). Thus, a discrepancy of at least a factor of \( \sim 3 \) is evident between our \( V \)-band \( \omega(\theta) \) measurements plotted in Figure 13 and the models plotted in RY’s Figure 3. Other observations that seem to rule out the most extreme merger galaxy evolution models (Carlberg 1995) are a collection of redshifts for very faint galaxies obtained with the Keck telescope (Koo et al. 1996). This spectroscopic sample is still sparse, so the results should be treated as preliminary, but the median redshifts obtained for \( I > 22 \) are contrary to what is expected for Carlberg’s “maximal merging model.” The extent of the role of galaxy merging is still not clear but we can conclude that a maximal merger model is no longer a viable mechanism for galaxy evolution at intermediate redshifts.

5.2. Color-selected Samples

With the multicolor deep imaging in each field, the angular correlation function can be measured with color-selected samples down to faint magnitude limits. A few approaches for obtaining color-selected samples were attempted to maximize the number of galaxies and thereby improve the accuracy of the \( \omega(\theta) \) measurements over a significant range in color. However, it should be noted that having just \( V, R, \) and \( I \) images to work with, and no bluer bandpasses, unfortunately leads to a limited baseline of observed colors for the faint galaxies. In Figure 15, the angular correlation functions selected by \( (V - R), (R - I), \) and \( (V - I) \) colors are presented in a plot analogous to Figure 1 of LSK. Amplitudes measured for the full range of angular separations, \( 10'-126' \) (yielding an effective separation of \( \sim 35'' \)), are plotted as a function of the colors that the galaxies observed are either less than (top row of plots), or greater than (bottom row). Poisson error bars are shown but should be considered to be underestimates of the true errors. The final amplitudes have been corrected for the integral constraints and stellar dilution factors calculated for each field. Also, the plotted values are obtained with weighted averaging, where the weights are determined from the number of galaxies detected in both bandpasses for a given color.

LSK used \( (U - R_g) \) colors to find that the \( \omega(\theta) \) amplitude increased by over a factor of 10 for the reddest and bluest galaxies taken from a bright magnitude-limited sample \( (B_g < 23.5) \). Figure 15 shows no indication of this behavior for the current, fainter galaxy sample with \( (V - R), (R - I), \) and \( (V - I) \) colors. Within the errors, the angular correlation amplitude integrated over the full range of separations is relatively constant regardless of the colors of the galaxies being analyzed. This is consistent with what BSM (\( R \approx 25.5 \)) and Infante & Pritchet (1995, \( b_j \leq 24 \) and \( R_p \leq 23 \)) have found, although one should note that these previous studies made only one division in color (blue/red) and did not look at the extremely blue or red galaxies, as did LSK. It is possible the lack of an increase in the clustering amplitude in Figure 15 is due to the \( V, R, \) and \( I \) colors not discriminating the reddest and bluest galaxies well enough with our coarser color bins. In other words, the numbers of galaxies available for this study may be simply insufficient for providing an accurate measure of \( \omega(\theta) \) with color. There also could be physical reasons for the nondetection of a clustering increase with extreme color, and these are discussed below.

To do a more direct comparison of the clustering observed with color-selected samples the correlation amplitudes for \( (V - I) \)-selected galaxy samples are plotted in Figure 16, for a fixed angular separation of \( 1' \). Results from NRGCI are given as open symbols for the 50% blue, 50% red, and entire samples, while the filled symbols are mea-
measurements for objects in the current work with \((V-I) < 1.3\), \((V-I) > 1.3\), and for the full sample, with the median \(I\) magnitude plotted on the abscissa. The value \((V-I) = 1.3\) was chosen as the dividing line such that the entire sample could be cut into roughly 50% blue and red galaxies. In this case, the \(\omega(\theta)\) amplitudes calculated from our sample assume that \(\delta = 0.7\) in order to ease the comparison with NRGCI, and our error bars are calculated using bootstrap resampling. Galaxies with the 50% reddest \((V-I)\) colors in NRGCI were observed to have clustering \(\sim 4-8\) times stronger than the blue half of the sample, but with substantial errors. NRGCI also argue that there is an increase in amplitude for the 20% color marginals (bluest and reddest \(\frac{1}{2}\)) of the galaxies from the 50% samples, although these two samples for either blue or red objects are consistent within the errors. The \((V-I)\)-selected sample in this study, which is 1 mag deeper than that of NRGCI, shows no sign of color segregation of the clustering amplitudes beyond \(I \sim 22\), albeit with large error bars for the measurements. Also, there is no significant difference observed between the clustering of the galaxies in the full, \(I\)-selected sample and the blue and red samples. The amplitudes from our red and blue galaxy samples do not bracket those calculated for the full sample due to not all the galaxies being detected in both the \(V\) and \(I\) images. Clustering amplitudes for the red and blue 20% marginals in our sample were not determined since large errors would result from the small sample size.

LP also determined \(\omega(\theta)\) for \((V-I)\)-selected samples of galaxies, in the magnitude range \(I = 18-20\). They used \((V-I) = 1.5\) as the blue/red boundary and found a marginally significant difference between the samples with red galaxies exhibiting stronger clustering, similar to the NRGCI results at brighter magnitudes. This comparison of LP, NRGCI, and the current work suggests that a difference in

the clustering amplitudes for blue and red galaxies [using \((V-I)\) selection] exists at bright magnitudes \((I \sim 18-21.5)\) and either disappears or has not been detected at fainter magnitude limits \((I \sim 21.5-24)\). Interpreting these color-selected clustering results is complicated in that, for \(I\)-selected samples (for e.g., the CFRS with \(I \leq 22\)), red galaxies tend to be confined to a fairly narrow range of intermediate redshifts, while blue galaxies are observed to have more broadly distributed redshifts \((z \sim 0-1)\), with a lower mean \(z\). The results summarized in Figure 16 may be showing that significant galaxy evolution is occurring at faint magnitudes relative to brighter magnitudes (lower redshift). Red and blue galaxies are observed to cluster differently at lower \(z\), which is simply a reflection of the morphology-density relation. Another possibility is that the blue sample is more diluted with lower luminosity galaxies that have stronger clustering properties, making the clustering measurements for the blue and red faint galaxy samples indistinguishable. A more accurate approach for tracing clustering evolution of “typical” \(I^\ast\) galaxies may be to select out red galaxies and measure their clustering variations with magnitude, since the luminosity function of these objects shows very little change over \(0 < z < 1\) (Lilly et al. 1995a). Obviously, larger multicolor imaging surveys of faint galaxies are required to more accurately measure the color-selected angular correlation function and further check the viability of various galaxy evolution scenarios.

![Figure 16](image)

**Fig. 16.—Amplitudes of the angular correlation function calculated for separations of \(r\) assuming \(\delta = 0.7\), following Fig. 3 of NRGCI to ease the comparison with this work. The abscissa is the median \(I\) magnitude for each galaxy sample plotted. Open symbols show the “50% Blue,” “50% Red”, and full samples of NRGCI. Filled symbols show the red, blue, and entire galaxy samples for this study, with \((V-I) = 1.3\) used as the blue/red dividing line.**

5.3. Comparison with Models of the Spatial Correlation Function

To determine a viable model for the evolution of \(\xi(r)\) with redshift (eq. [7]) from measurements of \(\omega(\theta)\), one requires the redshift distribution \((dN/dz)\) of the galaxies within the magnitude interval being considered. With a realistic redshift distribution, Limber’s equation (8) can be solved and relationships between the observed \(\omega(\theta)\) and inferred \(\xi(r)\) can be determined for different cosmologies \((q_0, H_0)\) and clustering evolution \((\epsilon)\), with the power-law index for \(\xi(r)\) constrained from the angular clustering results \((\gamma = \delta + 1)\). Unless otherwise noted we assume that \(q_0 = 0.5\) and \(H_0 = 100\) km s\(^{-1}\) Mpc\(^{-1}\). Since the correlation length at \(z = 0\), \(r_\text{c}\), corresponds to the amplitude of the locally observed \(\xi(r)\), a range of reasonable values for \(r_\text{c}\), and \(\epsilon\) are assumed in order to generate models of the clustering evolution. These models are then compared to the observed values of \(\omega(\theta)\).

The \(I\) photometric data in this study have a magnitude limit \((I \sim 24)\), which is a full 2 magnitudes fainter than the currently largest deep redshift survey (Lilly et al. 1995b). For an estimate of the redshift distribution at the limits of the photometry, the \(dN/dz\) measured to \(I \sim 22\) can be extrapolated to fainter magnitude limits with a no-evolution assumption for the galaxies. Evolution is obviously occurring for the galaxies at some level toward fainter magnitudes but the discrepancy between the observed galaxy counts and extrapolated \(dN/dz\) number counts is fairly small for the 2 mag interval beyond \(I \sim 22\) (see Figs. 8 and 9 in Lilly et al. 1995a). Using the Lilly et al. extrapolations to \(I \sim 24\) and the observed redshift distributions for brighter magnitude limits, we have calculated the variation of \(\omega(\theta)\) with \(I\) magnitude for given values of \(r_\text{c}\), \(\epsilon\), and \(\gamma\).

Errors are almost certainly present in the extrapolated redshift distributions used to calculate the clustering evolution models. Since the amplitude of \(\omega(\theta)\) calculated using Limber’s equation (eq. [8]) has a strong dependence on the
shape (essentially the width) of \(dN/dz\), errors will occur if this shape is poorly estimated with the extrapolation, while the effect of an inaccurate normalization will be small. Preliminary observational support for the extrapolated Lilly et al. (1995a) redshift distributions has been provided by the DEEP survey (Koo et al. 1996) where the median redshifts measured at faint limits are found to be consistent, albeit within large error bounds. Hudon (1995) and HL have shown that if the redshift distribution, with slightly brighter magnitude limits than ours, has 15% more galaxies added to it that are similarly distributed in redshift, there is little change in the estimate of the spatial correlation function, as expected. In the more extreme scenario where this 15% is added at median redshifts of \(z_{\text{med}} = 1.5\) or 2.1 as a Gaussian distribution the resulting correlation length \(r_0\) is increased by \(\sim 15\%\) or \(\sim 30\\%\), respectively. This is probably a reasonable upper bound for the uncertainty in this study due to our still sparse knowledge of the redshift distribution of galaxies with \(22 < I < 24\). Also, the shape of \(dN/dz\) at these faint magnitudes will be incorrect if a particular galaxy population dominates at these limits but is not detected at brighter magnitudes. The calculations using Limber’s equation are presented keeping these caveats in mind.

LP have measured \(\omega(\theta)\) for a wide range of brighter magnitudes in \(I\) and we use these results as a comparison to this study, as well as to the models. In Figure 17 the logarithm of \(\omega(\theta)\) at an angular separation of \(1\) plotted with the median \(I\) magnitude given on the abscissa. All of the samples obtained from the two fields (CL and FBS) in LP are included, in addition to the EBKTG point for the \(I\) band. Models calculated using the aforementioned redshift distributions from Lilly et al. (1995a) are plotted as a series of lines for different values of \(r_0(z = 0)\) (solid lines, \(5.4 \ h^{-1} \ Mpc\); dashed lines, \(4 \ h^{-1} \ Mpc\); dotted lines, \(2 \ h^{-1} \ Mpc\)) and \(\epsilon = -1.2, 0, 1, 2\) from top to bottom for each set of lines with a given \(r_0\). The value of \(\gamma\) is fixed at 1.8 following the discussions in §§4.2 and 4.6, but the effects of varying it are shown below. The LP results are amplitudes obtained from galaxies within narrow luminosity bins (1 or 2 mag wide) spanning \(I = 16 – 23\), while the points determined in the current study are for galaxies with \(I = 19 – 23\), 23.5, 24. Our magnitude limits were chosen to minimize the error in the \(\omega(\theta)\) measurements since there are a limited number of galaxies available (2697 with \(I = 19 – 24\)). As noted earlier, the majority of the galaxies in our \(I\)-selected samples are within a narrow magnitude range so no claim is made for a detection of flattening in \(\omega(\theta)\). When combined with the LP clustering study, a generally smooth decline in the amplitude of \(\omega(\theta)\) with \(I\) magnitude is observed. For the slightly better statistics of the \(R\) observations (Fig. 14), a decrease in the clustering amplitudes with magnitude limit is unambiguously seen to the faintest limits, in conjunction with other studies. There is fairly good agreement between the correlation function amplitudes from the three \(I\)-band studies at the faintest magnitudes in Figure 17. At \(I_{\text{med}} \sim 22\) the LP data are consistent with our measurement of the clustering. For the LP point at \(I_{\text{med}} \sim 22.5\) and the Efsta-thiou et al. result at \(I_{\text{med}} \sim 23\) there is agreement within 2\(\sigma\) of the amplitudes obtained from this study but with a substantial error for the LP measurement at their magnitude limit.

A comparison of the observations to the models in Figure 17 leads to some general conclusions. Clustering evolution that is fixed in comoving coordinates \((\epsilon = -1.2)\) is a viable scenario only if \(r_0(z = 0) \sim 2 - 3 \ h^{-1} \ Mpc\). Values for \(r_0\) are typically not observed to be this small for the entire galaxy population. The correlation length usually ranges from \(\sim 4 \ h^{-1} \ Mpc\), calculated using \textit{IRAS}-selected redshift surveys (Saunders, Rowan-Robinson, & Lawrence 1992; Fisher et al. 1994), to the canonical optical survey correlation length of \(r_0(z = 0) = 5.4 \ h^{-1} \ Mpc\) from Davis & Peebles (1983). BSM show that, with a correlation length of \(r_0 \sim 2 \ h^{-1} \ Mpc\), and a rate of clustering growth predicted by linear theory \((\epsilon \sim 0.8)\), they can match their clustering observations and models at faint limits. From this result, BSM claim that low surface brightness (LSB) and/or dwarf galaxies are dominating the faint galaxy population since some local measurements of the correlation lengths for these objects yield \(r_0 \sim 2.3 - 2.7 \ h^{-1} \ Mpc\) (Santiago & da Costa 1990). However, the values for the LSB/dwarf galaxy local correlation lengths are still controversial and may be larger (Thuan et al. 1991). As noted earlier (§5.1), the amplitudes of the BSM observations in the \(R\) band are significantly lower than what is observed in this study at similar magnitude limits. Given that most studies to date have found local correlation lengths with \(r_0 \gtrsim 4 \ h^{-1} \ Mpc\) along with the assumption that faint galaxy populations evolve into locally observed galaxies, our \(I\)-filter observations then suggest that \(\epsilon \geq 0\), in agreement with HL, Le Fèvre et al. (1996) and Shepherd et al. (1997).

For a non-negative value of \(\epsilon\), two general possibilities remain for the evolution of the faint galaxy population. The first scenario is that \(\epsilon \simeq 0 – 1\) and \(r_0(z = 0) \sim 4 \ h^{-1} \ Mpc\), where the excess of faint blue galaxies is due to objects that are analogous to \textit{IRAS}-selected galaxies with respect to star formation, morphology and clustering, as suggested by

![Figure 17](image-url)
BTBJ. The second possibility is that $\epsilon > 1$, implying significant evolution in the clustering from faint limits to locally observed galaxies such that a value of $\sim 5-6$ $h^{-1}$ Mpc is found for $r_0$. This correlation length is in agreement with most optically selected, local redshift survey measurements of $\xi(\theta)$. We note that EBKKG only considered clustering models with $-1.2 < \epsilon \leq 0$ in order to obey the standard gravitational instability picture. However, more recent $N$-body studies (Melott 1992; Yoshii et al. 1993) have found that models with $\epsilon \sim 0-3$ are indeed possible due to the continual merging of groups as the universe expands.

To illustrate the sensitivity of the models for $\omega(\theta)$ in $I$ to the assumed $q_0$, $\gamma$ and $\epsilon$, for a given $r_0$, the observed clustering amplitudes are plotted again in Figure 18 versus the $I$ magnitude limit along with three different families of models. Note that each ordinate of the three panels covers a different range of clustering amplitudes but has a total range of 1 dex. For each $r_0$ listed (5.4, 4.2 $h^{-1}$ Mpc from top to bottom) in the lower left corner, the solid line corresponds to the model calculated for $q_0 = 0.5$, $\gamma = 1.8$ and $\epsilon = 0$. Assuming a small-$q_0$ universe with $q_0 = 0.1$ yields the dotted line for each $r_0$. Changing just the power-law index for the correlation function to the two extremes of what is observed, $\gamma = 1.9$ and $\gamma = 1.6$, gives the short-long-dashed lines above and below the solid line, respectively. Finally, the long-dashed lines are associated with changing just the value of $\epsilon$, as in Figure 17. The long-dashed line above the solid line is for $\epsilon = -1.2$, while the two dashed lines below correspond to $\epsilon = 1, 2$ for decreasing amplitude. Clearly, the comparison between the observations and models does not have a strong dependence on the assumed value of $q_0$. The clustering evolution parameter $\epsilon$ provides the greatest leverage in parameter space for matching the observations to clustering models, in addition to being poorly constrained. If the inherent degeneracy of fitting clustering models to measurements of $\omega(\theta)$ is to be broken, better determinations of $r_0$ and $\gamma$ for local samples of galaxies selected by morphology, luminosity, and surface brightness are required.

6. DISCUSSION

Lilly (1993) summarized the three standard scenarios that were suggested to explain the obvious evolution in the galaxy number counts (Tyson 1988; Lilly et al. 1991) and the lack thereof in the measured redshift distributions (Broadhurst, Ellis, & Shanks 1988; Colless et al. 1990, 1993). First, the faint blue galaxies are explained to be protodwarf galaxies undergoing bursts of star formation at intermediate redshifts ($z \sim 0.4$) and then evolving into galaxies at the faint end of the luminosity function ($L < 0.1L^*$) by the current epoch (Broadhurst et al. 1988). The second model has faint galaxies being very short-lived, star-bursting objects that are subsequently disrupted or fade away in such a fashion that they are not observed in large numbers at small redshift (Babul & Rees 1992). The final conventional model invokes merging of subgalactic units at intermediate redshifts where current $L^*$ galaxies are the products of this process (Broadhurst, Ellis, & Glazebrook 1992). We refer to these three frameworks as the “bursting dwarfs,” “fading dwarfs,” and “merger” models. By looking for increases in the clustering of faint galaxies using various selection criteria (e.g., magnitude-limited samples, small angular separations, colors), we can test the viability of the merger galaxy evolution model, as in § 5.1. Dwarf models can be checked by comparing the clustering behavior of the various local and faint galaxy populations (§§ 5.2 and 5.3). A more unconventional explanation for the number counts and redshift distributions evolution discrepancy is the proposal by McGaugh (1994, also see Ferguson & McGaugh 1995) of the existence of a significant population of low surface brightness galaxies that are not typically detected locally, due to observational selection effects, but that can be detected in faint, photometric surveys.

As shown in the previous section, neither a sustained flattening in the amplitude of the angular correlation function over a significant range of magnitude, nor a well-defined change in the slope of the decreasing $\omega(\theta)$ with magnitude, is seen with current data, which suggests that merger models (Guiderdoni & Rocca-Volmerange 1990; Broadhurst et al. 1992; Carlberg & Charlot 1992) may not be viable descriptions of faint galaxy evolution. It is still possible to incorporate some merging into galaxy evolution models without being inconsistent with the observed clustering and redshift distributions, but models with “maximal merging” are certainly ruled out (§ 5.1 and Koo et al. 1996). A lack of a significant amount of merging of galaxies at faint limits is in agreement with our results from Woods et al. (1995), where no substantial excess of close pairs of galaxies in NF1 were found down to $I \leq 24$. Obviously, the clustering measurements displayed in Figure 14 are still too inaccurate to reasonably constrain any detailed galaxy evolution model.

Assessing whether the faint galaxy population are predominantly bursting dwarfs or fading dwarfs is very difficult (if not impossible) to distinguish at this juncture using only
estimates of the clustering. These two scenarios can be respectively described as follows (see Fig. 17 and Table 8): (1) the majority of faint galaxies are evolving into a local population similar to IRAS-selected galaxies ($r_o \sim 4 h^{-1}$ Mpc) with moderate clustering evolution ($\epsilon \sim 0 – 1$) and (2) the clustering is fixed in comoving coordinates with $\epsilon \sim -1.2$ requiring the local counterparts of faint galaxies to be either low surface brightness galaxies that are weakly clustered ($r_o \sim 2 – 3 h^{-1}$ Mpc) or galaxies that have been disrupted and are therefore undetectable. Additional local surveys for low surface brightness galaxies are required to better constrain the purported weak clustering behavior of this population. The third model, previously discussed, where merging of galaxies plays some role is given by (3) clustering of the faint population evolving at a significant rate ($\epsilon > 1$) yielding local galaxies with similar clustering properties to what is observed in optical surveys ($r_o \sim 5 – 6 h^{-1}$ Mpc). The three approximate pairs of $r_o$ and $\epsilon$, which match the clustering model to our $I$-band $\omega(\theta)$ measurements, are also in broad agreement with those found by HL in the $R$ bandpass. We summarize these three galaxy evolution scenarios in Table 8 and comment on the models’ strengths and drawbacks.

Clustering evolution processes (1) and (3) could both be occurring but recent HST observations suggest that late-type and irregular galaxies dominate the number counts at faint limits (Driver et al. 1995; Glazebrook et al. 1995). Loveday et al. (1995) have measured the local correlation lengths for early and late-type galaxy morphologies to be $5.9 \pm 0.7 h^{-1}$ Mpc (E and S0) and $4.4 \pm 0.1 h^{-1}$ Mpc (Sp and Irr), respectively. If spirals and irregulars are the dominant population comprising the faint counts it is possible that the primary path for faint galaxy evolution is scenario (1) above, with a moderate change in the clustering ($\epsilon \sim 0 – 1$). This value of $\epsilon$ is consistent with what is predicted by linear theory and is easier to account for than larger values of $\epsilon$.

Another point that should be made is that some of the large fraction of irregular galaxies found at faint magnitude limits with HST data (Driver et al. 1995; Glazebrook et al. 1995) could be galaxies at high redshift with $2.3 \leq z \leq 3.5$ (Abraham et al. 1996), where the lower redshift limit corresponds to the Lyman discontinuity entering the $U$ filter. This possibility is supported by 19 of the 83 “irregular/peculiar/merger” objects found by Abraham et al. in the Hubble Deep Field to have $(U – B) > -0.2$. A red UV-optical, $(U_v – \Phi)$, color criterion has also been recently used by Steidel et al. (1996) to find “normal,” star-forming galaxies (with $M_I \leq 25$) at redshifts of $3 \leq z \leq 3.5$, confirmed with Keck spectroscopy. However, Steidel et al. have found the surface density of the galaxies in this redshift range to be quite low ($0.4 \pm 0.07$ galaxies arcmin$^{-2}$). This suggests that if the Abraham et al. irregular objects are at high redshift then the bulk of these galaxies will have redshifts with $2.3 \leq z \leq 3$. Galaxy morphologies at large redshifts are subject to $K$-corrections, evolutionary effects of stellar populations and a strong dependence on surface brightness (Giavalisco et al. 1996). Considering these effects, it is not clear whether the significant fraction of faint galaxies with apparent irregular/peculiar morphology are all the result of dynamical evolution at intermediate redshifts or the galaxies are at high redshift ($z > 2.3$) and are observed to be irregular due to these “morphological K-corrections.” Morphology studies of local galaxies at UV wavelengths and further faint galaxy spectroscopy (e.g., Koo et al. 1996) should address this concern.

No evidence was found in this study for a dependence of clustering on galaxy color (Figs. 15 and 16), such as what LSK found using $U – R_p$ colors at brighter magnitude limits. Nevertheless, Roche et al. (1996) find a difference in the clustering of red and blue galaxies at $B \sim 25.5$. They suggest that the observed clustering is consistent with the LSK result at $B \leq 23.5$. No significant differences in the clustering measured for red and blue samples of galaxies were found by EBKTG and BSM. NRGC, as mentioned earlier, find only marginal differences in the clustering for $(V – I)$-selected samples of red and blue galaxies at faint limits ($I > 21.5$), while larger discrepancies are observed for brighter magnitudes ($I < 21.5$). This roughly consistent clustering of red and blue galaxies at faint magnitudes is in agreement with what is observed for the $(V – I)$-selected samples of this work (Fig. 16). It remains to be seen if the difference (or consistency) of the clustering between red and blue galaxies is a function of the sample magnitude limit. The accuracy of the measurements of the color-selected $\omega(\theta)$ in the current work does not warrant further analysis. Larger photometric surveys of faint galaxies will be essential toward achieving more precise determinations of $\omega(\theta)$ for magnitude-limited and color-selected galaxy samples.

7. SUMMARY

The amplitude of $\omega(\theta)$ is found to decrease with magnitude limit when our $R$- and $I$-band results are combined with clustering determined by other authors. This observed monotonic decrease with magnitude rules out “maximal merger” galaxy evolution models. Angular correlation function estimates (in $I$) in the current study and LP were compared to galaxy clustering evolution models, generated with CFRS redshift distributions that were extrapolated to the faint limits of the photometry. The observed clustering of the faint galaxies can be explained with local correlation lengths for this population of $\sim 4$ or $\sim 5 – 6 h^{-1}$ Mpc for moderate ($\epsilon \sim 0 – 1$) or strong ($\epsilon > 1$) clustering evolution, respectively. Clustering evolution fixed in comoving coordinates ($\epsilon = -1.2$) is possible but requires smaller correlation lengths ($\sim 2 – 3 h^{-1}$ Mpc) than what is usually observed for local galaxies. No evidence is found for variations in clustering that are dependent on galaxy color in this study. Larger photometric surveys are required to confirm the stronger clustering amplitudes for faint, red galaxies found by Roche et al. (1996) and to improve upon the accuracy of

| Model          | $r_o$ ($h^{-1}$ Mpc) | $\epsilon$ | Comments                                           |
|----------------|----------------------|-------------|----------------------------------------------------|
| (1) Bursting dwarfs…….. | $\sim 4$         | $\sim 0 – 1$ | Acceptable at all limits; only moderate clustering evolution required |
| (2) Fading dwarfs ……..  | $\sim 2$          | $\sim -1.2$  | Untestable locally (?)                              |
| (3) Merging ……………….. | $\sim 5$          | $\sim 1 – 2$ | Excessive (unobserved) clustering; evolution may be required |

TABLE 8

SUMMARY OF GALAXY EVOLUTION MODELS
angular correlation function measurements at faint magnitude limits.

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