SMOOTHED PARTICLE HYDRODYNAMICS SIMULATIONS OF COUNTERROTATING DISK FORMATION IN SPIRAL GALAXIES

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ABSTRACT

We present the results of smoothed particle hydrodynamics (SPH) simulations of the formation of a massive counterrotating disk in a spiral galaxy. The current study revisits and extends (with SPH) previous work carried out with sticky particle gas dynamics, in which adiabatic gas infall and a retrograde gas-rich dwarf merger were tested as the two most likely processes for producing such a counterrotating disk. We report on experiments with a cold primary similar to our Galaxy, as well as a hot, compact primary modeled after NGC 4138. We have also conducted numerical experiments with varying amounts of prograde gas in the primary disk and an alternative infall model (a spherical shell with retrograde angular momentum). The structure of the resulting counterrotating disks is dramatically different with SPH. The disks we produce are considerably thinner than the primary disks and those produced with sticky particles. The timescales for counterrotating disk formation are shorter with SPH, because the gas loses kinetic energy and angular momentum more rapidly. Spiral structure is evident in most of the disks, but an exponential radial profile is not a natural by-product of these processes. The infalling gas shells that we tested produce counterrotating bulges and rings rather than disks. The presence of a considerable amount of preexisting prograde gas in the primary causes, at least in the absence of star formation, a rapid inflow of gas to the center and a subsequent hole in the counterrotating disk. For a normal counterrotating disk to form, there must be either little or no preexisting prograde gas in the primary, or its dissipative influence must be offset by significant star formation activity. The latter scenario, along with the associated feedback to the interstellar medium, may be necessary to produce a counterrotating disk similar in scale length and scale height to the primary disk. In general, our SPH experiments yield stronger evidence to suggest that the accretion of massive counterrotating disks drives the evolution of the host galaxies toward earlier (S0/Sa) Hubble types.

Subject headings: galaxies: evolution — galaxies: interactions — galaxies: kinematics and dynamics — galaxies: spiral — galaxies: structure — hydrodynamics

1. INTRODUCTION

There are only a handful of known cases of massive counterrotating disks in spiral galaxies to date, and yet counterrotation in spirals cannot be deemed a rare phenomenon by any standards. There are hints that it may be quite common, in fact, in early-type spirals, particularly S0s (Kuijken, Fisher, & Merrifield 1996). The origin of any counterrotating mass (gas or stars) within a spiral disk is an important unsolved problem with profound implications for the formation and evolution of all spiral galaxies, but the existence of a significant retrograde mass component (comprising anywhere from ~10%–50% of the total mass of the disk system) is a particularly intriguing question that threatens to radically alter our view of the evolution of spiral galaxies. The rogues’ gallery of spirals with such massive counterrotating disks currently boasts the following members: NGC 4550 (Rubin, Graham, & Kenney 1992), NGC 7217 (Merrifield & Kuijken 1994), NGC 4826 (Braun, Walterbos, & Kennicutt 1992), NGC 3626 (Ciri, Bettoni, & Galletta 1995), NGC 3593 (Bertola et al. 1996), and NGC 4138 (Jore, Broeils, & Haynes 1996).

Apart from the challenge they present to the traditional view of spiral galaxies that has evolved over the last few decades, counterrotating disks raise several questions about other astrophysical processes, such as the role and fate of gas in galaxy interactions, star formation in galaxies that contain counterrotating gas, the accretion rates and star formation histories of spiral galaxies in general, and the impact of counterrotating populations on the overall stability of the disk system. Even though a recent survey of S0s (Kuijken et al. 1996) found counterrotating gas in almost a quarter of the sample, none of these galaxies have counterrotating stars. Why is the counterrotating gas not forming stars? If star formation is inhibited in counterrotating disks, how does one explain NGC 4550 and others with stellar counterrotating disks?

It is very unlikely that counterrotating systems can be produced indigenously or as a by-product of the galaxy formation process. The theory of formation of a spiral galaxy from a spinning protogalactic cloud does not admit the possibility of bidirectional spin being imparted to the disk system. Subsequent accretion or merger events are a much more plausible explanation, and even these are severely constrained by the observed coldness of the counterrotating galaxies. Dissipationless mergers, especially between progenitors with comparable masses, can be ruled out in the general case, although Pfenniger (1997) has recently been able to produce a remnant resembling NGC 4550 with a collisionless merger of two spirals with special

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initial conditions. This leaves minor, gas-rich mergers and gas accretion or infall as the most promising candidates.

The most puzzling aspect of massive counterrotating disks in spiral galaxies is that the host galaxies appear quite normal in every other respect, and there is no evidence of excessive thickening of the primordial disks due to the accretion of the counterrotating disk. This suggests that the accretion process must not be a rapid or violent one, but there may be deeper implications here for the interaction histories of all spiral galaxies. A recent study of the tidal thickening of galaxy disks (Reshetnikov & Combes 1996) indicates that the ratio of scale length to scale height, $h/z_0$, is 1.5–2 times lower for interacting disks. But if any galaxy can be assumed to have undergone an interaction in the past, the thinness of the noninteracting sample proves that this ratio returns to its higher value after a certain amount of time (order of 1 Gyr). If it can be proved that a spiral can double its disk mass without destroying the primordial disk in the process, then the present-day appearance of spiral disks can no longer preclude such interactions in their past.

The claim that the thinness of spiral disks places a stringent limit on past accretion (Toth & Ostriker 1992) is further challenged by recent simulations showing that the halo absorbs a good portion of the orbital energy and angular momentum of the satellite in spiral-dwarf mergers (Walker, Mihos, & Hernquist 1996; Huang & Carlberg 1997) and by analytical results indicating that an isothermal halo may even shield the disk from an external tidal field due to a satellite (Murali & Tremaine 1997). A study of the lopsidedness of the disks of field spirals suggests that the accretion rate of spiral galaxies may be as high as one small ($\sim 10\%$ mass) companion every 4 Gyr ( Zaritsky & Rix 1997). This is still a very uncertain estimate, and a more accurate estimate of the accretion rate, and a knowledge of how much accretion a disk galaxy can withstand, are important questions from a cosmological point of view. Observations of counterrotating disks combined with a good understanding of how they form provide a striking new way to independently constrain these estimates, because the accreted matter can be easily distinguished in a counterrotating galaxy.

With the aim of understanding the origin of massive counterrotating disks in spirals and S0s, we have developed a numerical code and run hydrodynamical simulations to investigate the processes that are most likely to produce such bizarre systems. We have combined an $N$-body gravity solver with a gas dynamics particle code for this purpose. To study the basic parameters of the processes involved, we first adopted a quick-and-dirty "sticky particle" gas dynamics approach that allowed us to test various scenarios with relatively small investments of CPU time. In this way we were able to test massive counterrotating disk formation in a fiducial cold primary (Thakar & Ryden 1996, hereafter TR) as well as model the formation of the recently discovered counterrotating disk in the early-type spiral NGC 4138 (Thakar et al. 1997, hereafter TRJB). For each type of primary modeled, we have tested two theories of origin: adiabatic (secular) gas infall and a gas-rich dwarf merger. In order to produce a disk with opposite spin, both of these processes require a retrograde orbit of accretion for the infalling gas or satellite galaxy with respect to the primary’s spin. Gas infall works well for both late and early-type primaries, but a dwarf merger, especially with a substantial amount of dissipationless matter in the dwarf, is not viable for a cold primary, because it plays havoc with the primary’s disk.

An accurate representation of the gas in astrophysical systems typically requires that the hydrodynamical conservation equations be solved and the effects of physical processes such as shocks and viscosity be included. This requires the gas to be modeled as a fluid, but a good compromise can be achieved with a particulate representation if each gas particle is smeared over a finite volume and its physical properties averaged or smoothed over that volume. This smoothed particle hydrodynamics (SPH) approach (Lucy 1977; Gingold & Monaghan 1977; Monaghan 1992) meshes well with an $N$-body particle code but is much more costly in terms of CPU resources than the sticky particle approach. We therefore reserved it for a more detailed and restricted look at the structure of the counterrotating disk formed by reexamining (with SPH) a carefully chosen subset of the simulations we presented in TR and TRJB. We present the results of these simulations here. Although most of the simulations are reruns of those presented before with sticky particles, a couple of new runs are also discussed. Our SPH code currently does not include star formation. We intend to incorporate this in our simulations in the future, but as we discuss in § 4 below, we do not expect it to have a profound impact on the structure of the counterrotating disks formed.

Our computational method is described in the following section. Results of simulations are presented in § 3, followed by discussion and conclusions in § 4 and § 5, respectively.

### 2. Computational Method

Our SPH code is an implementation in the C language of the TREESPH code (Hernquist & Katz 1989, hereafter HK). We started with a basic C version of the tree code (Barnes & Hut 1986) that performs the force calculation, and optimized it for execution on the Cray Y-MP at the Ohio Supercomputer Center (see TR for details). These changes included rewriting the tree construction and traversal routines in such a way that they made extensive use of the vector capabilities of the Y-MP (Hernquist 1990; Makino 1990). The outcome was to speed up the execution by a factor of 5 or more. We then added the SPH code to it according to the prescription given by HK, using the same vectorization techniques as before to optimize the tree traversals required for nearest neighbor searches. Vectorization of the SPH calculations was tricky, and we benefited considerably from a perusal of the vectorized Fortran TREESPH code (a copy was kindly provided by Lars Hernquist). To test the final result, we ran the one- and three-dimensional tests described in HK to ensure that it gave identical results.

For the simulations described here, we have assumed an isothermal gas with a cutoff temperature of $10^4$ K. This has been shown to be a fairly good assumption for the interstellar medium (ISM) in disk galaxies, where most of the gas hovers close to the cutoff temperature because of the short timescale of the radiative cooling process (Barnes & Hernquist 1991; Mihos & Hernquist 1994a, 1994c). The motivation for assuming an isothermal gas is that it yields a tremendous saving in the CPU resources required for each simulation, since solving the energy equation is a nontrivial exercise that requires a semi-implicit scheme (see HK) and must be performed at the smallest timescale involved in the simulation.
Even with the isothermal assumption, and with the maximum possible vectorization, our SPH code typically requires 5–10 times the number of CPU resources required for the sticky particle code. This is mostly because the gas particles have their own individual time steps (each particle’s time step is a power of 2 subdivision of the system time step, so that all gas particles are synchronized at the end of a system time step, as discussed in HK), and this causes the performance to degrade when there are even a few particles in the lowest time step bin. To keep this problem under control, we do impose a minimum smoothing length (usually about a quarter of the softening length), but even so the isothermal assumption is more of a necessity than an option for us. As we discuss below, this assumption does present some problems when we attempt to model a gas-rich dwarf galaxy.

We have adopted the spherically symmetric spline kernel for SPH averaging (Monaghan & Lattanzio 1985; eq. [2.8] in HK). However, we do not use it for gravity softening. Our gravity solver uses the standard softening scheme, in which the gravitational potential is of the form \( \Phi \propto (r^2 + \epsilon^2)^{-1/2} \) (corresponding to a Plummer density profile), with softening length \( \epsilon = 0.5 \) kpc for all simulations. Although our code provides the capability to use different softening lengths for the collisionless and gas particles, we have used the same length for both species in the simulations reported here. The softening length effectively limits our resolution, but it is also necessary to suppress the two-body effects in our moderate-N simulations.

Our time-integration scheme is different from the one used by HK. Their scheme requires manual synchronization of the particle positions and velocities by applying midpoint corrections (eqs. [2.42]–[2.45] of HK). We decided instead to adopt a slightly different scheme that does not require these corrections (Katz, Weinberg, & Hernquist 1996). Although this scheme is not time-symmetric as a result of the particle positions and velocities being advanced in a different way than in HK, it is simpler and faster without introducing any additional error.

For all simulations reported here, we use the combined form of the bulk and Von Neumann-Richtmyer artificial viscosity, given by equations (2.24)–(2.25) in HK, and applied the restriction necessary to avoid viscous cooling of receding particles, as specified by HK. This viscosity introduces less shear than other artificial viscosities, although it does not provide as accurate a description of the flow near shocks. We felt that it was more important to avoid excessive shear in disk galaxy simulations and therefore adopted the combined form of viscosity.

We use 32K particles each to represent the primary’s halo and disk and up to 20K particles for the gas, depending on its mass. Where a dwarf galaxy is involved, the stars and dark matter in the dwarf are represented by enough particles so that the per-particle mass is roughly the same as in the case of the primary. A nagging problem with disk simulations is the disk heating due to the graininess of the potential that is inherent in a particle code (Hockney & Eastwood 1989). The problem is more severe in our simulations because of the longer timescales involved in retrograde interactions. Even with a softening length as high as 0.5 kpc, we find that there is considerable disk heating over the duration of our simulations with the number of particles we have chosen, and the heating of the disk due to the interaction itself is always of the order of the numerical heating. Unfortunately, the two are not separable, and this limits our ability to estimate the disk thickening accurately. Although we would like to use more than 32K particles each to represent the disk and halo, our CPU resource limitations currently prevent us from running considerably larger N-simulations.

The primary galaxy physical parameters are the same as we used before for the cold primary (TR) as well as for NGC 4138, the hot primary (TRJB). For convenience, they are repeated in Table 1. Our fiducial cold primary parameters are typical for an Sbc spiral like the Milky Way, and the disk is bar-unstable, as is consistent with its coldness (Binney & Tremaine 1987). Unless explicitly stated, the primary disk in most of the simulations reported here contains no primordial gas. The halo is a truncated isothermal sphere (truncated at \( r = R_h \)) with a logarithmic profile and core radius \( R_c \) as per equation (2-54a) in Binney & Tremaine (1987). Although most of the SPH simulations are reruns of previous sticky particle runs, there are some changes and some new simulations. The input parameters for the gas infall and merger simulations are listed in Tables 2 and 3, respectively. Correspondences to the original simulations in TR and TRJB are pointed out, where applicable, when we discuss the results in the following section. To avoid confusion with the sticky particle runs, we have named the SPH gas infall runs G1–G8 and the SPH dwarf merger runs D1 and D2.

The models for gas infall and mergers are the same as in TRJB. The infalling gas is configured as a long rectangular column of uniform density and square cross section. The gas columns we use in the SPH simulations may be shorter than the corresponding sticky particle simulations, primarily because the SPH runs require much longer to run and hence a shorter column with a correspondingly shorter infall time is preferable. The primary disk plane is selected to be the \( xy \)-plane, and the long axis of the infalling gas slab is parallel to the \( y \)-axis. The primary disk rotation is counterclockwise, and the center of the primary disk is at the origin of the coordinate system. The gas is then placed in the first quadrant in the same plane as the primary disk and is given an initial kick, mostly in the negative \( y \)-direction (an initial speed in the positive \( x \)-direction is also given in order to increase the initial angular momentum of the gas in some cases, see Table 2). This ensures that

### Table 1

| Parameters for Primary Galaxy Models |
|-------------------------------------|
| **Disk** | **Halo** |
| \( M_d \) (\( 10^{10} \times M_\odot \)) | \( M_h \) (\( 10^{11} \times M_\odot \)) |
| \( R_p \) (kpc) | \( R_h \) (kpc) |
| \( R_q \) (kpc) | \( R_c \) (kpc) |
| \( h_p \) (kpc) | \( h_c \) (kpc) |

| Model                  | \( M_d \) | \( R_p \) | \( R_q \) | \( h_p \) | \( M_h \) | \( R_h \) | \( R_c \) |
|------------------------|----------|----------|----------|----------|----------|----------|----------|
| Cold Primary (TR)      | 5.5      | 21.0     | 3.50     | 0.3      | 2.2      | 50.0     | 1.0      |
| NGC 4138 (TRJB)       | 2.0      | 6.0      | 1.25     | 0.2      | 1.5      | 20.0     | 5.0      |
the gas is accreted on a retrograde orbit with respect to the primary disk spin.

For the dwarf mergers, the setup is basically the same, except that the dwarf galaxy replaces the gas column in the first quadrant. The dwarf galaxy model is the same as described in TR and TRJB. We have only attempted two gas-rich dwarf mergers (D1 and D2), both with NGC 4138 as the primary, since they were not feasible for the cold primary because of the problems discussed in TR. Both mergers have identical input parameters and are a repetition of merger M2 from TRJB. SPH is used from the beginning for D1, whereas for D2 the merger is run with sticky particles until the dwarf galaxy is completely disrupted by the primary, at which point SPH replaces sticky particles for further calculations. This is because of the difficulties we encountered in using isothermal SPH to model the dwarf galaxy (see § 3.2.2).

3. RESULTS

The primary galaxy is evolved in isolation for a few dynamical times before the gas or dwarf galaxy is introduced. Hence the simulation results are shown starting at $t \gtrsim 1.5$ Gyr for the cold primary and $t \gtrsim 0.5$ Gyr for NGC 4138 (which has a smaller dynamical time). The simulation stops once most of the gas has settled in the plane of the primary and the evolution of the gas disk has tapered off to nearly zero.

3.1. Fiducial Cold Primary

For the cold primary, our first infall simulation is a repetition of the second continuous infall simulation discussed in TR, but in the SPH run we use a shorter column of gas. The infall simulation shown in Figure 1 has all other inputs the same as before (see Table 2 for input parameter values). A counterrotating disk with spiral structure has started to form by $t \sim 4.5$ Gyr. The side view confirms the disk formation and indicates that the counterrotating disk is uniformly thin throughout. Since the width of the last two panels in Figure 1 is 40 kpc, the size of the counterrotating disk appears to be comparable to that of the primary (radius 21 kpc). One has to be careful in making this comparison, however, since the radial profile of the counterrotating disk is not close to that of the primary. Saying that the sizes are comparable does not mean that the scale lengths are comparable. The lack of a centrally concentrated or exponential radial profile is clear from both the side and top views for $t \gtrsim 5$ Gyr. There is a hole in the inner regions of the counterrotating disk, and this, along with the larger size, could be the result of the gas having too much angular momentum initially.

Although the observed instances of counterrotating disks do not show a consistent radial profile, it is nonetheless important to determine whether a counterrotating disk with a similar scale length and profile as the primary disk can be formed. There is at least one example of this, NGC 4550 (Rix et al. 1992). We have experimented with lower values of the initial angular momentum of the infalling gas to determine what effect it has on the radial profile and structure of the counterrotating disk formed. Figure 2 presents the results of run G2, which has lower initial angular momentum of the gas as well as a narrower (and hence denser) column of infalling gas than G1. Since the panel widths in Figure 2 are the same as in Figure 1, the differences are easy to see. The size and structure of the counterrotating disk formed are quite different. The disk is smaller (by about 25%–30%) than the disk formed in G1, and the radial profile is more centrally concentrated. There is a lot more structure visible, since the gas has piled up in several places because of shocks. The side view shows that the disk is as thin as in G1, and it is significantly thinner than the primary disk.

A comparison of the primary disk at the beginning and the end of G2 with the counterrotating disk is shown in Figure 3. The primary disk shows a bar even before the gas is introduced, but the bar does not become more pronounced after the formation of the counterrotating disk. It is, in fact, obscured by the inclination of the disk in the top view of the final time step. The final thickness of the primary disk is hard to measure because of its inclination, but we do measure it by compensating for the inclination and computing the $z = 0$ plane at each location in the disk (see below). However, it is obvious from Figure 3 that the primary disk thickness has more than doubled, and it now resembles a lenticular (SO) disk rather than an Sb disk. This is further corroborated by the rotation curves and velocity dispersion plots for the final disks shown in § 4, where we also discuss some of the possible reasons for the pronounced heating experienced by the cold primary.

### Table 2

| Run  | $M_\text{gas}^a$ | $V_\text{gas}^b$ | $V_\text{ro}^c$ | $V_\text{vo}^d$ | $X_\text{0}^e$ | $L_\text{0}^f$ | $T_\text{0}^g$ |
|------|------------------|------------------|------------------|------------------|------------------|------------------|------------------|
| G1   | 4.40             | 0.3              | 0.10             | 50               | 50               | 300              | 20               |
| G2   | 4.40             | 0.3              | 0.00             | 50               | 50               | 300              | 10               |
| G3   | 1.32             | 0.3              | 0.05             | 25               | 25               | 150              | 10               |
| G4   | 1.32             | 0.3              | 0.05             | 25               | 5                | 100              | 10               |
| G5   | 1.45$^b$         | 0.3              | 0.10             | 25               | 10               | 100              | 10               |
| G6   | 1.76$^b$         | 0.3              | 0.05             | 25               | 5                | 100              | 10               |
| G7   | 1.06$^i$         | 0.2              | ...              | ...              | 30               | ...              | 10               |
| G8   | 1.06$^i$         | 0.1              | ...              | ...              | 30               | ...              | 10               |

$a$ Total gas mass ($10^{10} M_\odot$).
$b$ Initial velocity of the gas in the negative $y$-direction, specified in units of the centripetal velocity [$G(M_\text{h} + M_\text{d})/X_\text{vo}^e$]^{1/2}, where $M_\text{h} + M_\text{d}$ is the total mass of the primary (halo plus disk).
$c$ Initial velocity of the gas in the positive $x$-direction, specified as a fraction of $V_\text{ro}^c$.
$d$ Initial $y$-distance of the gas column from the center of the primary disk (kpc).
$e$ Initial $x$-distance of the gas column from the center of the primary disk (kpc).
$f$ Length of the infalling gas column (kpc).
$g$ Thickness of the gas column (side of square cross section [kpc]).

### Table 3

| Run       | $M_\text{g}^a$ | $M_\text{s}^b$ | $M_\text{d}^c$ | $E_\text{s}^d$ | $R_\text{s}^e$ | $R_\text{d}^f$ | $R_\text{c}^g$ |
|-----------|----------------|----------------|----------------|---------------|---------------|---------------|---------------|
| D1/D2     | 1.06           | 0.26           | 2.64           | 0.27          | 4.0           | 4.0           | 8.0           |

$a$ Mass of gas ($10^{10} M_\odot$).
$b$ Mass of stars ($10^{10} M_\odot$).
$c$ Mass of dark matter ($10^{10} M_\odot$).
$d$ Ratio of gas mass to total mass.
$e$ Radius of stellar sphere (kpc).
$f$ Radius of gas sphere (kpc).
$g$ Radius of halo (kpc).
The counterrotating disk is of comparable size but is less than half as thick as the primary disk. The velocity fields of the primary disk and the counterrotating disk are shown in Figure 4. The primary's velocity field is considerably hotter than the velocity field of the counterrotating disk, but they are clearly antiparallel.

The radial profiles of the primary disk and the counterrotating disks at the end of simulations G1 and G2 are compared with each other in Figure 5. The dip in the particle density between ~1–3 kpc is clearly visible for G1, and although there is a slight dip at those radii even for G2, the G2 profile comes much closer to the exponential radial profile of the primary disk.

The thickness of the primary disk is plotted for simulations G1 and G2 in Figure 6. A comparison of the thickness plots with those in TR shows some similarities. The thickness of the postaccretion primary disk in TR was only slightly higher than the thickness of the isolated primary disk evolved over the same time period. The thicknesses for G1 and G2, however, are significantly higher (by about
25%) than the thickness plotted for continuous infall in TR. This is probably because of the fact that we are using a smaller softening length (0.5 kpc instead of 1.0 kpc used by TR), even though we have doubled the number of particles used to represent the halo for the simulations in this paper (we used 16K particles for the halo in TR). The thickness curve for the isolated disk is also higher by a similar amount compared to the corresponding curve in TR. Some of the increase in the heating of the primary can also be attributed to the fact that the infalling gas column is shorter, and hence more dense, than the one used in TR, and the timescale for formation of the counterrotating disk has decreased by \( \geq 30\% \) (see § 4.1). This gives the primary disk less time to adjust. Furthermore, the smaller softening length aggravates the gravitational impact of the incoming gas on the primary.

We are not so concerned with the heating of the primary in the simulations reported here, because we have established in TR and TRJB that if the rate of accretion is low enough, the effect on the primary can be kept to an acceptable level. As mentioned above, our motivation for hastening the counterrotating disk formation process by using a
shorter column of gas than before was to save CPU time expended on each simulation. The fact that this causes considerably more heating of the primary underlines the sensitivity of the cold primary to the rate of accretion.

3.2. Hot Primary: NGC 4138

3.2.1. Gas Infall

The first infall simulation for NGC 4138, G3, is a repetition of simulation I4 from TRJB, although the total mass of the gas is 25% higher in the SPH runs. The remaining parameters are the same as in the sticky particle run. The results are shown in Figure 7. The counterrotating disk has started to form at $t \sim 1.8$ Gyr, and after $t \sim 2.6$ Gyr, it does not evolve very much. The size of the counterrotating disk is significantly larger than the primary, since it extends over more than two-thirds of the width of the panel (20 kpc) in the side view at $t \geq 2.6$ Gyr, but it clearly does not have an
exponential radial profile. There is a small central mass concentration. The lack of an exponential radial profile is easier to see in the side view at $t = 3.0$ Gyr, because the disk is slightly tilted to the line of sight, and an outer ring is visible in addition to the central concentration, with a drop in the particle density at intermediate radii. The radial density distribution is plotted below in Figure 10. As in the case of the cold primary (G1), the lack of an exponential profile indicates that a good fraction of the particles do not lose enough angular momentum to allow them to settle in the inner few kpcs of the disk. Spiral structure is evident, but it is not very strong. The side view at $t = 2.6$ Gyr shows the counterrotating disk to be quite thin, although this is hard to see in the last panel because of the inclination of the disk.

The velocity fields of the primary and counterrotating disk are compared in Figure 8. The larger size of the counterrotating disk is also evident in the velocity field plot. The velocity field of the counterrotating disk is colder than that of the primary and more regular.

A lower angular momentum version of G3 is attempted in G4, shown in Figure 9. The initial velocity of the gas column is the same as G3, but the initial distance of the gas from the center of the primary is smaller by a factor of 5 (see Table 2). The length of the gas column is also two-thirds of that in G3. The parallels between G1/G2 and G3/G4 are quite obvious. Once again, the lower angular momentum produces a counterrotating disk with a more centrally concentrated radial profile and a smaller overall size. A comparison of the radial profiles of G3 and G4 with each other and the primary’s radial profile is shown in Figure 10. Even though the radial profile of the counterrotating disk in G4 is not very close to the exponential profile of the primary, it is much closer to that than G3. There is no pronounced dip in the particle density in the inner radii for G4, as there is for G3.

There is not much difference in the heating (and thickening) of the primary disk with the SPH simulations of NGC 4138 compared to TRJB. Although the heating is a little higher on average in our SPH simulations, this is because of the fact that we are using shorter gas columns. This makes the accretion rate higher. In the cases where a lot of gas does make it to the nuclear portions of the primary, there is slightly more heating of the primary. In the other cases, the levels are comparable to those obtained in TRJB (after allowing for the shorter gas columns), and hence the thickness plots are not repeated here.

### 3.2.2. Gas-Rich Dwarf Merger

Our previous studies have established that the gas-rich dwarf merger model of counterrotating disk formation is only feasible for a hot primary like NGC 4138. TR found that even a merger with a 10% mass satellite produced unacceptable levels of heating in a cold primary disk, whereas the hot primary used by TRJB fared much better in this regard and was able to withstand a gas-rich dwarf merger under restricted conditions.

Apart from the issue of whether a dwarf merger can produce a counterrotating disk, we encountered a technical difficulty with our gas-rich dwarf galaxy model. A stable model of a collisionally supported gas-rich dwarf galaxy is not easy to obtain with an isothermal equation of state, because the gas in the dwarf tends to shock and collapse rapidly if the temperature is too low and expand if it is too high. This inability of the isothermal model to yield a stable configuration is not very surprising, since isothermal pressure gradients do not grow rapidly enough to combat the gravitational collapse of the gas (Tohline 1980). A rotationally supported dwarf would probably not alleviate this situation very much, because it would still collapse in the direction not affected by the centrifugal force, thereby yielding a quasi-equilibrium isothermal disk. As such, we restrict ourselves to testing only one model for which we have a reasonable expectation of success, and even for this we have to adjust the gas temperature to keep the dwarf galaxy in equilibrium.

We have selected merger M2 from TRJB, with a dwarf that is $\sim 18\%$ as massive as the primary. Two versions of this merger are presented here. In the first case (run D1), we use SPH with a higher isothermal temperature until the dwarf galaxy is tidally disrupted by the primary, and thereafter reduce the temperature of the gas to the usual value ($10^4$ K). In the second version (run D2), we use the sticky particle results to evolve the dwarf until it forms a thick disk around the primary and then switch to SPH with the usual settings.

The results for D1, with SPH being used all the way, are shown in Figure 11. The top view shows that the dwarf makes one entire pass around the primary before getting completely stripped and ends up forming a very barred counterrotating disk. In the side view, the disk appears to be quite thin, and the velocity field diagrams shown in Figure 12 confirm that a disk is indeed formed. The stellar particles (not shown) of the dwarf form a thick flattened cloud around the primary, in the same way as they did with sticky particles. The velocity fields also indicate that the primary disk is heated up more than in the infall case (compare with Fig. 8).

When SPH is introduced at a late stage after the dwarf has already been tidally stripped and has formed a thick disk around the primary, the evolution of the disk is very different (Fig. 13). D2 dramatically illustrates the differences between sticky particle gas dynamics and SPH. Within a short time ($\lesssim 0.4$ Gyr), the thick sticky particle disk has collapsed to a thin disk. There is a slow accumulation of
mass in the center thereafter, but beyond the formation and dissipation of a few rings, the disk is fairly stable.

It appears that the nature of the dissipation is very different with SPH. With sticky particles, there is a slow but steady loss of kinetic energy, whereas with SPH, shocking causes rapid loss of energy. The isothermal model probably contributes to the speed of the energy loss, since there is no counteracting increase in thermal pressure to oppose the kinetic energy dissipation until the cutoff temperature is reached. Once the SPH gas forms a kinematically cold disk, evolution ceases as the dissipation drops off suddenly. The rapid loss of kinetic energy is demonstrated even more effectively when prograde gas is present in the primary (see below) and the counterrotating gas collides head-on with it.

3.3. Other Experiments with NGC 4138
3.3.1. Prograde Gas in the Primary

Sticky particle simulation I6 (TRJB), which included a ring of prograde gas in the primary, showed that the prograde gas did not have much of an impact on the kinematics of the counterrotating gas. Collisions between counter-
streaming gas particles did send some gas to the center, but this was a modest effect and the prograde gas ring remained mostly intact.

The SPH simulation G5, shown in Figures 14 and 15, is a repetition of I6 from TRJB. The average radius of the ring is 1.7 kpc and the width of the ring is 1 kpc (the inner radius of ring = 1.2 kpc, and the outer radius = 2.2 kpc). The mass of the ring is \sim 10% of the mass of the infalling gas, and the ring has uniform density. The particles in the ring are given initial velocities appropriate for primary disk particles at those radii, and the ring is checked for stability for a couple of dynamical times prior to introducing the infalling gas.

We find that the behavior of the gas in the SPH version of this simulation is dramatically different from the sticky particle results. Virtually all of the prograde gas interacts with the counterrotating gas and falls to the center, leaving a gap in the counterrotating gas distribution at the original location of the prograde ring of gas. The interaction between the prograde and retrograde gas particles can be seen more clearly in the velocity fields plot. The angular momentum of the retrograde gas is reversed near the center of the primary as both the prograde and retrograde gas particles fall to the center.

Going one step further, we included a prograde disk of gas in the primary along with the stellar disk. The mass of the gas disk is \sim 20% of the mass of the primary disk, and the radius of the gas disk is 4 kpc. As was the case for the prograde ring, the prograde disk is of uniform density, and we evolved it with the rest of the primary first for 200 Myr (without any counterrotating gas) to ensure that it was stable on its own. Then we introduced the infalling gas at \( t = 0.7 \) Gyr. This is simulation G6, shown in Figure 16. The behavior of the two gas components is similar to that seen in G5, only stronger. The prograde gas sweeps up all the retrograde gas that it comes in contact with, and both combine to form a nuclear mass concentration within 1 Gyr or so. As in the case of the prograde ring, the velocity field plots (Fig. 17) show that the angular momentum of the infalling gas is reversed in the process. The rest of the infalling gas that does not come into contact with any of the prograde gas forms a counterrotating outer ring. The gap between the prograde nuclear disk/bulge and the outer counterrotating disk is larger in this case, since the prograde gas occupied a larger area in G6 than in G5.

### 3.3.2. Infalling Spherical Gas Shell

Gas infall is not well constrained by observations, and, as an alternative gas infall model, we tried a spherical shell of gas that is given retrograde angular momentum with respect to the primary disk. The two runs with such an infalling gas shell, G7 and G8, have the input parameters shown in Table 2. G8 has half the angular momentum of G7. The results for G7 and G8 are shown in Figures 18 and 19. The shell collapses to a flat disk within a very short time (\sim 0.5 Gyr). The speed of the collapse is somewhat higher for G8. The final size of the disk in G8 is about half the size of the disk in G7 (note that the last two panels of Fig. 19 are half as wide as the corresponding panels in Fig. 18). The mass distribution of the two counterrotating disks is also different. There is an outer ring present in both disks, but only G8 shows a significant central mass concentration. The tilting of the disk at the end of G8 is in phase with the tilting of the primary disk. The formation of the outer ring in both cases, together with the dependence of the size of the ring on the initial angular momentum, suggests that the infalling shell mechanism is not particularly suitable for counterrotating disk formation. There is no a priori reason to believe that an infalling gas shell is physically more realistic, and it certainly does not offer any advantages over the infalling gas column that we have chosen for most of our infall experiments.

The formation of rings in the collapse of the spherical gas shells is not a big surprise. Ring formation in the collapse of rotating spherical systems has been documented by several workers, with transient rings forming in dissipationless collapses (Miller & Smith 1979) and more stable rings forming in the collapse of rotating isothermal gas clouds (Tohline 1980; Boss & Haber 1982). The equilibrium of self-gravitating isothermal gas rings has been investigated from a theoretical perspective as well (Ostriker 1964). Unlike the
The radial distribution of the gas is much closer to an exponential disk, rather than a ring as in G3.

The rings we see are not transient structures that are formed en route to a more permanent morphology. Furthermore, they are true rings in that they are not due to orbit crowding, and the particles in them are not constantly moving in and out of the ring configuration. Our rings are consistent with those seen and studied in isothermal gas cloud collapse scenarios and are self-gravitating: the mass per unit length is much greater than the equilibrium mass per unit length for isothermal rings (eq. [156] in Ostriker 1964). For G7, this quantity is approximately four times, and for G8, it is approximately seven times the equilibrium value at the last time step, which corresponds to $\sim 11t_{\text{ff}}$ for G7 and $\sim 8t_{\text{ff}}$ for G8 (where $t_{\text{ff}}$ is the free-fall time of the initial spherical shell). However, collapse is most likely inhibited, because our numerical resolution (as set by the softening length, which is 0.5 kpc) is of the order of the ring width. Our values of $\alpha$ and $\beta$, the ratios of the thermal and rotational energies to the gravitational energy, respectively (Boss & Haber 1982), are 0.02 and 0.12 for G7 and 0.02 and 0.07 for G8. The value of $\alpha$ is well below the maximum value that will produce a collapse. For G7, the $\beta$ value falls within—whereas for G8 it falls below—the 0.1–0.5 range seen by Boss & Haber (1982) for isothermal rings. However, a lower
produced a counterrotating disk in $\gtrsim 6$ Gyr (TR), whereas G1 and G2 take $\gtrsim 4$ Gyr. For NGC 4138, sticky particle simulations were able to produce a counterrotating disk in $\gtrsim 3$ Gyr (TRJB), whereas G3 and G4 achieve the same result in $\gtrsim 2$ Gyr. Most of the decrease in the time required is due to the shorter columns of gas used in the SPH simulations, but a comparison of the evolution of the sticky particle (we ran a sticky particle simulation with the same inputs for comparison) and SPH gas also shows that the latter loses its kinetic energy and angular momentum more rapidly. We estimate that the difference in the timescales due to the gas dynamics alone is of the order of 5%–10%. The infalling gas is captured and assimilated faster into the primary disk because of the more efficient viscous dissipation in SPH.

Strangely enough, the infalling gas does not collapse and fragment like it did with sticky particles. The basic character of dissipation appears to be different: in sticky particle dissipation, head-on collisions were not quite so dissipative, and there was a comparable amount of dissipation regardless of how much collisional support the gas had; in SPH, whenever collisions between gas particles cause shocks, the dissipation is drastic and the gas quickly loses most of its collisional kinetic energy. As a result, as long as the SPH gas motion does not involve any head-on collisions between opposing gas streams, the gas dissipates very little energy. The tendency of SPH to cause rapid loss of kinetic energy and angular momentum in counterstreaming gas has been noted by other users of SPH as well (J. C. Mihos 1997, private communication), and it is possible that SPH is too dissipative in nature (the artificial viscosity is too high) under these circumstances.

4. DISCUSSION

4.1. SPH versus Sticky Particles

Our primary aim with sticky particle simulations (TR; TRJB) was to determine the range of input parameters that worked and to investigate the effect of the infall/merger on the primary disk. We succeeded in achieving these objectives, and we did not put much faith in the sticky particle results where the structure of the counterrotating disk was concerned. The SPH simulations presented here confirm that we were justified in our mistrust of the sticky particle results. Shocks and viscosity paint quite a different picture of counterrotating disk structure.

The most visible differences in the disks formed with SPH are their thinness and radial structure. Sizes are comparable to the sticky particle disks, although the SPH disks do on average tend to be a little smaller. However, they are all significantly thinner than their sticky particle counterparts. We found the sticky particle disks to be on average thicker than the primary disks. With SPH the reverse is true. The velocity field diagrams also show that all the counterrotating disks are considerably colder than the primary disks. This was true even with the sticky particle disks.

There is more structure visible in the SPH disks. A comparison of the radial density profiles of the counterrotating disks for runs G3 (dotted line) and G4 (dashed line), measured at the final time steps. The surface particle number density $N$ is calculated as in Fig. 5 and plotted as a function of radius $R$. The radial profile of the primary disk at the end of G3 (the G4 primary profile is very similar) is also plotted for reference (solid line).

The timescale of the counterrotating disk formation is significantly smaller (by $\gtrsim 30\%$) with SPH. A comparison of simulations G1-G4 with their sticky particle counterparts proves this. The sticky particle run corresponding to G1

\[ \beta \text{ limit for isothermal rings has not been established as yet, and it should be remembered that our initial configuration is a spherical shell rather than a cloud.} \]

4.2. Source of Infalling Material

While it is fine to theorize about the formation of counterrotating disks by gas infall or even gas-rich dwarf infall, the evidence for such material in the intergalactic medium is still quite thin, which is the main reason why there are no realistic models of infall available. The possibility of undetected molecular gas in the intergalactic medium around spirals provides one glimmer of hope. There has been more evidence recently for a significant fraction of the neutral ISM in our galaxy being H$_2$-bearing molecular clouds (Dixon, Hurwitz, & Bowyer 1998). The general case for ongoing gas infall in spiral galaxies is also getting stronger, with recent evidence from IRAS data (albeit still somewhat controversial) that massive star formation rates are independent of spiral type and that the median gas recycling time is $\sim 1 \times 10^8$ yr (Devereux & Hameed 1997), pointing to a sustainable outside source of gas.

"Galaxy harassment" has been recently suggested as a means of spiral evolution in clusters (Moore et al. 1996). Galaxy harassment refers to multiple, high-speed encounters that transform small spirals into dE/dSph, but more importantly, leave giant tidal debris arcs and tails that could provide fuel for quasars as well as the raw material for producing counterrotating disks. The recent discovery of a giant debris arc in the Coma cluster (Trentham & Mobasher 1997) supports this theory. Harassment of the debris tails themselves would create tidal shocks, leading to formation of dwarf galaxies in the tidal tails, as shown also in numerical simulations (Barnes & Hernquist 1992). This would mean that the incidence of counterrotating disks in
clusters should be higher than in the field, something that can be tested for in the future as more systematic surveys of counterrotation in clusters (especially the more distant ones) are undertaken.

The main difficulty posed by dwarf mergers is the heating experienced by the primary disk due to the large amounts of dark matter currently believed to be associated with most dwarf galaxies. There is still some doubt whether the evidence for the dark matter in dwarf galaxies is trustworthy or not, and if vindicated, dwarf spheroidal satellites without dark matter (Kroupa 1997) may provide a better chance of obtaining counterrotating disks from dwarf mergers. The instances of counterrotating bulges in spiral galaxies (e.g., NGC 7331; Prada et al. 1996) and counterrotating cores in ellipticals certainly point to merger events (Balcells & Quinn 1990). We are able to create a counterrotating bulge with our gas-rich dwarf model when the dwarf is sufficiently dense and massive (simulations S1 and S2 in TR), but the dissipationless (dark and stellar) matter in the dwarf is a severe liability to the survival of the primary disk.

4.3. Gravitational Influence of Various Components

To compare the gravitational influence of the three components (halo, disk, and gas) in the disk plane, we have
plotted the gravitational force experienced by a test particle of unit mass in the midplane of the disk ($z = 0$) as a function of radius in Figure 21. This is shown at the beginning and end of the simulation, for simulations G2 (cold primary) and G4 (hot primary). The gravitational forces experienced by the test particle due to the halo particles, the primary disk particles, and the gas particles, are shown separately. The forces are further separated into their azimuthally averaged radial (in plane of disk) and vertical components. Values for the outer halves of the disks are less reliable because of low particle numbers, and for G2 the calculation of the force components is further complicated by the substantial inclination of the primary and gas disks at the end of the simulation. We have attempted to correct for this inclination as best we can by first applying a coordinate transformation to reset the disk plane to the $xy$-plane before doing the force calculations.

For the cold primary (G2), initially ($t = 1.5$ Gyr) the primary disk exerts the dominant radial force in the inner half of the disk system. The disk and halo radial forces are almost equal in the outer half. The vertical forces due to all components are initially an order of magnitude lower than the radial forces. As expected, the influence of the infalling gas is initially negligible. This is because the gas is several disk radii away at the beginning of each simulation. At the end of the simulation ($t = 6.0$ Gyr), the halo-disk force distribution with radius does not change significantly, but the magnitude of the forces is lower on average by $\sim 7\%$.
radial force due to the gas is now comparable to the halo and disk radial forces. In the first couple of kpc, the gas radial force dominates. The vertical influence of the halo has increased and that of the primary disk has diminished by a considerable amount (~50%–60%), especially in the outer half of the disk plane. The gas now exerts a vertical force that is dominant in the inner half of the disk and is intermediate between the halo and disk vertical forces in the outer half. On average, though, the vertical forces are still an order of magnitude lower than the radial forces.

For the hot primary (G4), the initial (t = 0.5 Gyr) breakdowns are similar, although the halo is even less dominant in the inner half of the disk plane because of its high core radius (5 kpc). The vertical forces are, as before, an order of magnitude lower than the radial forces, and the gas influence is initially negligible. At the end of the simulation (t = 2.5 Gyr), the influence of the gas is close to that of the halo in the inner regions but is significantly less than the disk and halo forces in the outer disk. The vertical influence of the halo remains much the same, whereas the vertical force due to the primary disk increases in the inner half. The gas vertical force rivals that due to the halo throughout and is not the dominant vertical influence, as was the case for the cold primary. The gas mass fraction (compared to the primary disk mass) is much less in the hot primary, and so it is not surprising that the gas has a smaller influence both radially and vertically.

4.4. The Nature of the Gas Disks

The final rotation curves and midplane volume densities of the counterrotating gas disks at the end of simulations G2 and G4 are shown in Figure 22. The rotation velocity and velocity dispersion values (azimuthally averaged) are shown in the plots at the top, with the SPH-computed total midplane gas density (azimuthally averaged) at each radial location shown by filled circles in the bottom plots. This midplane density is compared to $\rho_{00}$, the unperturbed central density for the plane-parallel approximation obtained by adding up the contributions due to the gas.
pressure and the (external) gravitational pressure, as per equation (9) in Elmegreen & Elmegreen (1978):

$$\rho_{00} = \left[ P_{\text{ext}} + (1/2)\pi G \sigma^2 \right]/c^2.$$

This "expected" midplane density is plotted as open circles, and the second term on the right of the above equation, due to the gas surface density, is plotted as open squares. The gravitational pressure in the first term is obtained from the combined halo, primary disk, and gas vertical forces shown in Figure 21.

For the cold primary (G2) gas disk, the rotation velocity is initially steep and levels off at $R \sim 3$ kpc, whereas the hot primary (G4) gas disk's rotation curve has a more gentle slope toward its flat portion. This is because of the higher halo core radius for the NGC 4138 model, which reduces the halo's influence in the inner few kpc. The heights of the flat portions of the rotation curves are comparable for the gas disks in G2 and G4, but the G2 gas disk clearly shows a higher average velocity dispersion. The counterrotating disk formed in the cold primary is evidently somewhat hotter than the one formed in the hot primary.

The value of $\rho_{00}$, the expected midplane density obtained by adding up the two contributions from the gravitational and gas pressures, compares well with the actual computed midplane gas density, although there are significant differences. For the cold primary, the expected value is consistently less than the actual density, except for the outer disk. In the case of the hot primary, the reverse is true. The differences seen in the computed and expected densities are probably due to errors in the surface density and gravitational pressure values obtained from the particle data. Sources of uncertainties include the cutoff height assumed for the gas disk, the correction applied for the disk center-of-mass and inclination (especially in G2), and lower particle numbers at higher radii. The curves for the gas pressure (surface density) term indicate that the gas pressure is the dominant contributor for the gas density for the hot primary, whereas for the cold primary the gravitational pressure is the dominant influence, especially in the outer disk, where the gas surface density falls off considerably.

4.5. Heating of Primary Disk

The initial and final rotation curves for the primary disks after gas infall are shown in Figure 23 for both the cold (G1–G2) and hot (G3–G4) primary simulations. The final rotation velocities and velocity dispersions are consistent with those observed for S0/Sa galaxies (e.g., Seifert & Scorza 1996; Rix et al. 1992), since $v_{\text{rot}}/\sigma \lesssim 2$ rather than $> 1$, as is
true for Sb and later types. Furthermore, it is clear from Figure 23 that both the cold and hot primaries end up looking like S0s after the formation of the counterrotating disk. The velocity dispersion ranges from $\sim 100$ to $\sim 150$ km s$^{-1}$ for the final disks in either case, with the rotation velocities dropping by $\sim 50$–$100$ km s$^{-1}$ for the cold primary and by $\sim 25$–$30$ km s$^{-1}$ for the hot primary. In other words, the cold primary fares much worse than the
primary that was hot (S0/Sa) to begin with. In the same
vein, the outer portions of the primaries, which are colder
initially (since the velocity dispersion is initially proportion-
al to the surface density), are heated up more than the inner
regions.

As mentioned above, the heating seen in simulations
G1–G4 is more than was seen for the corresponding sticky
particle simulations, because the infalling gas columns are
shorter in the SPH simulations. However, the fact that
colder disks are more susceptible to heating is independent
of the actual amount of heating experienced. Although this
vulnerability of cold disks is hardly a surprise, it does
strengthen the argument for predicting that most instances
of massive counterrotating disks are likely to be found in
S0/Sa galaxies. This is consistent with the observations to
date.

4.6. Tilting of Disks

It is apparent from most of the simulations presented
above that the resulting counterrotating gas disk is signifi-
cantly tilted with respect to the initial symmetry plane of the
primary disk (the $z = 0$ plane). This is not a phenomenon
new to our SPH simulations. It was observed in the sticky
particle simulations as well, and we have commented on it
previously (TRJB). It is worth reiterating here that the
tilting experienced by the gas disk is in phase with the tilting
of the primary stellar disk, it occurs over relatively short
timescales (order of dynamical time of primary disk) com-
pared with the overall timescale of each simulation, and it is
in response to the combined torques exerted by the halo
and the gas.

Even though the net gravitational torques and the
angular momentum vectors are initially all aligned with one
another, the forces experienced by individual particles sub-
sequently are by no means all in the initial plane of the
primary disk. This is especially true for the gas particles,
which undergo gas dynamical interactions. For instance in
G2, which shows the maximum final disk inclination, the
side views for $t = 3.0$ and $t = 4.0$ Gyr (Fig. 2) show an
excess of gas particles below the disk plane (i.e., in the lower
half of the panels, which are centered on $z = 0$) on the left
side of each panel. This excess mass concentration below
the plane on one side is consistent with the subsequent tilt
of the gas disk (which tilts in phase with the primary disk)
seen in later panels.

Comparisons of the total torques exerted by the halo and
the infalling gas on the primary disk with the rate of change
of the angular momentum vector (TRJB) of the disk show
that the two quantities are consistent with each other. The change in the value of the z-component of the total angular momentum vector, which is by far the dominating component, is less than 1% over the course of the simulation in most cases and up to ~3% in the extreme cases.

4.7. Star Formation

Our SPH code does not yet incorporate star formation. Previous studies suggest that in interactions of galaxies where substantial amounts of gas are involved, the star formation rates are very modest for more than 90% of the interaction history, with almost all of the star formation occurring in a rapid starburst after the gas has dissipated most of its kinetic energy and formed dense regions in the centers of the galaxies (Mihos, Richstone, & Bothun 1992; Mihos, Bothun, & Richstone 1993; Mihos & Hernquist 1994a, 1994b). The prestarburst phase of the star formation in these simulations had no discernible effect on the final structure of the gas, since the depletion rates were very low. In our case, the longer timescales involved in our retrograde simulations may deplete significantly more gas over the course of the simulation, but other than reducing the intensity of the central starburst, this is unlikely to have a profound effect on the structure of the gas disks formed as long as there is little or no prograde gas in the primary disk. If there is a considerable amount of preexisting prograde gas, then it is possible that star formation resulting from the head-on collision of the counterstreaming gas particles will prevent a significant amount of gas from falling to the center because of rapid conversion to stars. This may alleviate the problem of excessive dissipation that causes the gas to collapse to the center of the primary in our simulations that include prograde gas (G5 and G6). The presence of star formation may also be instrumental in producing counterrotating radial profiles that are closer to the primary radial profiles (i.e., closer to exponential profiles). We plan to include star formation in our code in the future, but for now we have excluded it to save precious CPU time.

5. CONCLUSIONS

There are some notable differences between the characteristics of the counterrotating disks resulting from SPH and those obtained with sticky particle simulations by TR and TRJB. The SPH disks are very thin compared to their sticky particle counterparts and to the primary disks. They also show evidence of spiral structure, and their size and
Fig. 20.—Top and side views of the counterrotating disk formed in G4 showing evidence of lopsidedness (top) and warps (bottom).

radial mass distribution are quite sensitive to the input parameters, particularly those that affect the initial angular momentum of the gas.

Other differences in the SPH results include the lack of clumping of the infalling gas, a problem that was quite severe with our sticky particle simulations, and the shorter timescales for disk formation.

Although it is easy to produce thin counterrotating disks with gas infall, it is not so easy to obtain exponential radial profiles. The initial angular momentum of the gas has to be low enough, and some combination of other processes, such as prograde gas, star formation, and energy feedback from massive stars, may be necessary to produce counterrotating disks that are very similar to the primary disks. Currently, there is no evidence to indicate that counterrotating disks have predominantly exponential profiles, so this is not necessarily a problem.

In general, the process that dumps a massive counterrotating disk in a cold primary spiral, especially if it is a minor gas-rich merger but even if it is gas infall that occurs over a few dynamical times ($\lesssim 10t_{\text{dyn}}$), is likely to heat up the primary substantially and change its type to an S0/Sp galaxy. On the other hand, if the primary is already an S0/Sp galaxy to begin with, then it can acquire a counterrotating disk without changing its type significantly. The fact that most of the currently known instances of massive spiral counterrotating disks are in S0/Sp galaxies is therefore a selection effect rather than an accident.

The presence of primordial prograde gas in the primary has a drastic effect on the retrograde gas that comes in contact with it. Neutralization of the angular momenta is rapid, with both the prograde and retrograde gas particles ending up in the center of the primary within a few dynamical times. This may be an indication of a problem with SPH that causes overdissipation in counterstreaming gas flows, at least in the absence of star formation. The inclusion of star formation and energy feedback from supernovae will most likely yield significantly different results in such situations.

A retrograde-rotating, infalling gas shell produces a counterrotating bulge and flat outer ring but is unable to produce a counterrotating disk in the proper sense. The size of the ring is well correlated with the angular momentum of the shell. The formation of the ring is consistent with previous studies of collapsing isothermal gas clouds. These studies also suggest that significantly hotter gas with lower angular momentum is necessary to produce a counterrotating disk with this model.

We hope to test our results further in the near future with the addition of thermal effects and star formation to our SPH code.

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Fig. 21.—Azimuthally averaged gravitational force due to the halo (circles), disk (squares), and gas (triangles) particles, at points located in the plane of the disk at the plotted radii, with the component in the plane shown as the solid line and the component perpendicular to the plane shown as the dotted line. The plots are shown for the initial (top panels) and final (bottom panels) configurations, with results for the cold primary simulation G2 being shown on the left and the hot primary simulation G4 shown on the right. The force values are in simulation units, and time is in Gyr.
Fig. 22.—Final rotation velocity (open circles) and velocity dispersion (open squares) plots (top panels) and the midplane gas density (bottom panels) for the gas disks formed at the end of simulations G2 (left panels) and G4 (right panels). In the gas density plots, the actual average midplane gas density computed using SPH is plotted as filled circles, whereas $\rho_{\text{tot}}$, the expected midplane density (see text), is shown as open circles. The contribution to the gas density due to the gas pressure (see text) is plotted as open squares. The density values are in simulation units.
Fig. 23.—Comparison of the rotation velocity $v_{\text{rot}}$ and velocity dispersion $\sigma$ as a function of cylindrical radius $R$ of the initial primary disks (top) with the final primary disks (middle and bottom) for gas infall simulations with a cold primary (a) and hot primary (b). The open circles show the mean rotation velocity, and the open squares show the velocity dispersion. The velocity unit is $10^3$ km s$^{-1}$. Both velocities for the final disks have been corrected for disk inclination, but values for $R < 0.1R_d$ where $R_d$ is the maximum value of $R$ plotted, are not reliable. The filled triangles show the ratio $v_{\text{rot}}/\sigma$ (limits shown on the right axis) as a function of radius.

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