

Theory of the Interaction of Planetary Nebulae with the Interstellar Medium

Ruth Dgani

*Department of Astronomy, The University of Texas at Austin, Austin
TX 78712*

Abstract. The theory of the interaction of planetary nebulae with the interstellar medium is important for the interpretation of nebular morphologies that deviate from point symmetry. It can be used to probe the interstellar medium and its magnetic field. We emphasize in this review the role of hydrodynamical instabilities in the interaction.

1. Introduction

The interaction of planetary nebulae (PNs) with the interstellar medium (ISM) was first discussed by Gurzadyan (1969). Smith (1976) used the thin shell model to analyze the evolution of a spherically symmetric expanding PN shell moving through the ISM (see also Isaacman 1979). On the observational side, only a few PNs showed interaction with the ISM (hereafter refereed to as interacting PNs) until the last decade. With better observational imaging instruments (CCD cameras) the field interacting PNs became more popular (e.g., Borkowski, Tsvetanov, & Harrington 1993; Jacoby & Van de Steene 1995; Tweedy, Martos & Noriega-Crespo 1995; Hollis *et al.* 1996; Zucker & Soker 1997). Borkowski, Sarazin & Soker (1990; hereafter BSS) discuss the potential in using interacting PNs to probe the ISM, and list many interacting and possibly interacting PNs. This list was extended in the review article by Tweedy (1995) and in the atlas of interacting PNs (Tweedy & Kwitter 1996). Tweedy with several collaborators has been leading the observational research in the field (e.g., Tweedy & Kwitter 1994a,b; Tweedy & Napiwotzki 1994). Other studies are in progress, and preliminary results were presented in the conference (e.g. Kerber *et al.* 2000; Dopita *et al.* 2000).

On the theoretical side, Soker, Borkowski & Sarazin (1991; hereafter SBS) performed numerical simulations, both in the isothermal and adiabatic limits. Their results validate the thin-shell approximation of Smith (1976). In the adiabatic case they found that the Rayleigh-Taylor (RT) instability plays a crucial role in the evolution. Recent numerical calculations reported in this conference (Dopita *et al.* 2000; Villaver *et al.* 2000) confirmed these results. In their observational analysis of the interacting PN IC 4593, Zucker & Soker (1993) suggested that the interaction process oscillates between the adiabatic and isothermal cases. An analytical study of this process, the so-called radiative shock overstability, was performed by Dgani & Soker (1994) for finite-sized objects (e.g., spherical nebulae). The radiative overstability was found to occur only

in large PNs, with radii of $R \gtrsim 10^{19} n_0^{-1} \text{cm}$, where n_0 is the total number density of the ISM (in cm^{-3}). The instability occurs only for large PNs because the transverse flow, here it is the flow around the nebula, stabilizes the radiative overstability (Dgani & Soker 1994; see also Stevens, Blondin & Pollock 1992, for a numerical study of a different situation). For an earlier review of the theory see Dgani(1995). Several recent theoretical studies were conducted since that review. Soker & Dgani (1997) studied analytically the interaction of PNs with a magnetized ISM. Dgani & Soker (1998) and Dgani(1998) discussed the role of instabilities in the interaction. New numerical studies are underway (Villaver *et al.* 2000; Dopita *et al.* 2000), preliminary results presented in the conference show that instabilities play an important role in the interaction.

2. Preliminary Considerations

2.1. Why is the interaction important?

The interaction of PNs with the ISM causes deviation from axisymmetry (or point symmetry). Several other processes that cause deviation from axisymmetry involving binaries were discussed by Soker and collaborators (e.g. Soker 1994; Soker 1996; Soker et al 1998).

It is important to understand as accurately as possible the ISM effects on the morphology and to distinguish them from other processes.

Interacting PNs are also a tool for studying the ISM. This role was emphasized in previous reviews (Tweedy 1995; Dgani 1995).

2.2. Detecting the interaction

The interaction can be divided into at least three phases: free expansion, deceleration and stripping off the central star. Initially, PNs expand freely because their densities are of orders of magnitude higher than the densities of their surrounding medium. Eventually, though, every nebula will reach a stage in which the pressure inside the nebula is of the same order as the pressure outside the nebula. A strong enough pressure wave will then propagate into the nebula and decelerate it. If the nebula moves supersonically, the dominant ISM pressure is the ram pressure. The critical density n_{crit} , at which the ISM and the nebula reach the same pressure is (BSS, Eq. 1):

$$n_{crit} = \left(\frac{v_* + v_e}{c} \right)^2 n_0, \quad (1)$$

where c is the isothermal sound speed in the nebula, $c \sim 10 \text{ km s}^{-1}$ in most cases, v_* is the velocity of the central star, v_e is the expansion velocity of the nebula and n_0 is the ISM density. The most difficult parameter to observe directly in the above formula is the ISM density, n_0 . The nebula at this early stage of the interaction is not distorted, but its leading edge is brighter. Measuring the value of the density at the leading edge, n_{crit} , and using the above formula yields an estimate for n_0 .

If we take the average value of $v_e = 20 \text{ km s}^{-1}$ (Weinberger 1989) and $v_* = 60 \text{ km s}^{-1}$ (Pottash 1984; BSS), then $n_{crit} = 60n_0$. As the ISM ram pressure

depends on the velocity squared, the interaction is much stronger for fast moving planetaries. For example, if $v_* = 150 \text{ km s}^{-1}$ then the critical density is about 300 times that of the ISM.

The electron number density of extended planetaries is about 100 cm^{-3} (Kaler 1983) for the low surface brightness PNs from the NGC catalog, and 10 cm^{-3} for the fainter PNs from the Abell (1966) catalog. The very large nebula S-216 has $n_e = 5 \text{ cm}^{-3}$. In principle, therefore, fast moving planetary nebulae can be detected in very low density environments $n_0 \simeq 0.01$.

3. Numerical Models

The first realistic models of moving PNs were calculated numerically by SBS. They performed hydrodynamic simulations of the PN-ISM interaction with the particle in cell (PIC) method in cylindrical symmetry. They calculated two types of models of thick nebular shells moving through the ISM: an adiabatic model which applies to fast nebulae moving in the galactic halo and an isothermal model relevant to average nebulae moving in the galactic plane.

3.1. Cooling in the ISM shock

A nebula moving supersonically in the ISM creates a strong bow shock before it. The cooling time of the shocked ISM gas depends on the ionization state of the gas before the shock. The following analytical formula for the cooling time is based on numerical calculations of a steady shock with ionization precursor. It is accurate within a factor of 2 for $v_s > 60 \text{ km s}^{-1}$ McKee (1987):

$$t_{cool} = 100 \left(\frac{v_s}{50 \text{ km/s}} \right)^3 n_0^{-1} \text{ yr.} \quad (2)$$

In steady state shocks with velocities of about 50 km s^{-1} , the cooling drops by a factor of 10 when the material is preionized (Raymond 1979). However, the cooling time is still shorter than the flow time for an average nebula moving in the galactic plane. Only when the relative velocity is high and the density is low, will an adiabatic flow appear.

3.2. The adiabatic model

SBS choose the parameters for the adiabatic model so that the cooling time is larger than the flow time. An instability appears near the axis while the shell is decelerating which is interpreted as a Rayleigh-Taylor instability in the decelerated shell. When the shocked ISM cannot cool, the deceleration of the dense nebular shell by the dilute shocked ISM is Rayleigh-Taylor unstable. The main morphological features of the adiabatic flow are the ISM shock front, the decelerated PN shell and the RT instability. The last manifests itself as a "bump" and a hole in the nebula along the upstream symmetry axis. The numerical scheme used by SBS, though, being of 2D cylindrically symmetric nature and limited spatial resolution, forces the fastest growing mode to be of large wavelength and near the symmetry axis. Similar instability modes are obtained in numerical simulations of other types of dense bodies moving through an ambient media (Brighenti & D'Ercole 1995 for Wolf-Rayet ring nebulae, and Jones, Hyesung &

Tregillis 1994 for dense gas clouds). Brighenti & D’Ercole (1995) present a test run where the numerical grid is Cartesian rather than cylindrical. They find similar fragmentation of the flow, but not along the symmetry axis.

3.3. *The isothermal model*

In the isothermal model, efficient cooling is assumed. The cooling is calculated in the following way: for every cell in which the temperature is more than 10^4K the temperature is set equal to 10^4K . The model parameters are typical for an average nebula moving in the galactic plane.

In this case the shocked region is thin and "arms" of denser ISM material follow the nebula. The Rayleigh-Taylor instability does not appear in this case. This is because the shocked ISM is cooled and its density is of the same order as the nebular density. The model agrees with the thin shell approximation of Smith.

3.4. *Recent numerical simulations.*

Numerical simulations show that several types of instabilities which develop on the interface of spherically expanding shells (or winds) moving with respect to the ISM can fragment the shells (e.g. Brighenti & D’Ercole 1995). The detailed 2D numerical simulations of Brighenti & D’Ercole (1995) nicely show how RT and KH instabilities on the interface of the wind and the ISM develop, and allow the ISM to penetrate into the wind-bubble.

New numerical studies of interacting PNs are underway (Villaver *et al.* 2000; Dopita *et al.* 2000). Preliminary results presented in the conference show that the RT instabilities can fragment the nebular shell and allow the ISM to penetrate to the inner parts of the nebulae in some cases.

4. The Role of the ISM Magnetic Field in Interacting PNs

The important role of the ISM magnetic field on shaping interacting PNs was first mentioned in the observational paper of Tweedy *et al* (1995) about the nebula Sh 216. The morphology of the nebula is dominated by a bow shock followed by several parallel elongated filaments. Tweedy *et al.* (1995) noted that the magnetic pressure of the ISM is negligible compared to its ram pressure for the assumed central star velocity ($v_* \sim 20 \text{ km s}^{-1}$). They have argued that the central star velocity must be very low i.e. $v_* \sim 5 \text{ km s}^{-1}$ in order to explain the magnetic shaping of the nebula. Dgani & Soker (1998; table 1) noted that too many interacting PNs (a dozen) show signs of magnetic shaping i.e. parallel elongated filaments. It is very unlikely that all of them have such abnormally low central star velocity.

4.1. *General features*

Soker & Dgani (1997) conduct a theoretical study of the processes involved when the ISM magnetic field is important in the interaction. In the case where the ISM is fully ionized, we define four characterizing velocities of the interaction process: the adiabatic sound speed $v_s = (\gamma kT/\mu m_H)$ and the Alfvén velocity $v_A = B_0/(4\pi\rho_0)^{1/2}$ of the ISM, the expansion velocity of the nebula v_e , and the relative velocity of the PN central star and the ISM v_* . The interesting cases,

with the magnetic field lines being at a large angle to the relative velocity direction, are:

1. $v_* \gg v_A \sim v_s \sim v_e$, and a rapid cooling behind the shock wave. Both the thermal and magnetic pressure increase substantially behind a strong shock. If radiative cooling is rapid, however, the magnetic pressure will eventually substantially exceed the thermal pressure, leading to several strong MHD instabilities around the nebula, and probably to magnetic field reconnection behind the nebula.
2. $v_* \gg v_A \sim v_s \sim v_e$, and negligible cooling behind the shock. The thermal pressure, which grows more than the magnetic pressure in a strong shock, will dominate behind the shock. Magnetic field reconnection is not likely to occur behind the nebula. This domain characterizes the interaction of the solar wind with the atmospheres of Venus and Mars (e.g., Phillips & McComas 1991).

In the ISM it is likely that $v_s \sim v_A$, and $v_e \simeq 10 \text{ km s}^{-1}$ will in most cases not exceed the ISM sound speed. The central star velocity $v_* \gg v_s$ for most cases. Case (1) is the magnetic parallel of the isothermal model (section 3.1) which is relevant for planetary nebulae moving in the galactic plane. Case (2) is relevant for galactic halo PNs..

4.2. The magnetic Rayleigh-Taylor instability for galactic plane PNs

The isothermal case (case 1) is the more typical for the majority of PNs, moving moderately fast ($v_* \sim 60 \text{ km/s}$) close to the galactic plain, where the ISM density is relatively high. Soker *et al.* (1991) find in their numerical simulations that the isothermal model shows no sign of the RT instability. Soker & Dgani (1997) show the ISM magnetic field, which was not incorporated in the numerical simulations of Soker *et al.* (1991), can lead to an interesting manifestation of the RT instability in the isothermal case (case 1). We assume that there is a magnetic field of intensity B_s in the shocked ISM, but not in the nebula, and that its field lines are parallel to the boundary separating the ISM and nebular material. The acceleration g (deceleration in the frame of the central star) of the nebular front (pointing downstream) is in the opposite direction to the density gradient. In the linear regime the growth rate of the RT instability, σ , with a wavenumber $k = 2\pi/\lambda$, is given by (e.g., Priest 1984; §7.5.2)

$$\sigma^2 = -gk \frac{\rho_n - \rho_I}{\rho_n + \rho_I} + \frac{B_s^2}{4\pi(\rho_n + \rho_I)} k^2 \cos^2 \theta, \quad (3)$$

where ρ_I and ρ_n are the ISM and nebular density, respectively, θ is the angle between the magnetic field lines and the wave-vector, and RT instability occurs when $\sigma^2 < 0$. The deceleration of the leading shell by the ISM ram pressure is given by

$$g \sim \pi R^2 \rho_0 v_*^2 / M_F, \quad (4)$$

where R is the nebular radius, and M_F is the mass being decelerated. Taking $R = 0.5\text{pc}$, $v_* = 60\text{ km s}^{-1}$, $\rho_0 = 10^{-25}\text{ g cm}^{-3}$ and $M_F = 0.05M_\odot$ we obtain $g \sim 3 \times 10^{-7}\text{ cm s}^{-2}$.

We assume that after it is shocked the nebular material reaches 10^4 K , and that its thermal pressure $\rho_n v_{sn}^2$ is approximately equal to the ram pressure of the ISM, $\rho_0 v_*^2$. Here ρ_n is the shocked nebular density, $v_{sn} \simeq 10\text{ km s}^{-1}$ is the isothermal sound speed in the shocked cooled nebula, and we use the assumption $v_* \gg v_e$. The post-shock nebular density is therefore

$$\rho_n = \rho_0 v_*^2 / v_{sn}^2. \quad (5)$$

RT instability occurs when the density of the shocked cooled ISM ρ_s is smaller than the post-shock nebular density. This can happen for a magnetic shock when the magnetic pressure provides the support for the shocked ISM region. For a fast enough growth of RT instability we require $\rho_s / \rho_n < 1/3$. Assuming equipartition in the undisturbed ISM, we find the condition for a strong enough RT. Namely, we find that for a typical PN velocity $v_* = 60\text{ km s}^{-1}$ through the ISM the pre-shocked ISM magnetic field line should be within 45° of the shock front direction. Therefore, in most typical cases, i.e., equipartition and $v_* \sim 60\text{ km s}^{-1}$, fast growing RT instability modes with long wavelengths will develop.

The magnetic field has a destabilizing effect only for modes having wave numbers close to being perpendicular to the field lines (and in the plane separating the ISM and the nebula). Instability modes having wave numbers along the field lines are suppressed by the magnetic tension. This effect may create elongated structures in the direction of the magnetic field lines.

4.3. The morphology of interacting PNs and their galactic latitude

Dgani & Soker (1998) explore the observational consequences of the development of RT in galactic plane PNs moving in a magnetized ISM. In order to define the galactic plane and galactic halo population we use the expression for the cooling time after the ISM shock (eq. 2). The cooling will be effective when the flow time of shocked ISM around the nebula $t_{flow} \sim R/v_*$ is longer than the cooling time. Calling the ratio of the two times η , we obtain:

$$\eta = t_{cool}/t_{flow} \sim 0.1 \left(\frac{v_*}{50\text{ km s}^{-1}} \right)^4 \left(\frac{R}{10^{18}\text{cm}} \right)^{-1} \left(\frac{n_0}{0.1\text{cm}^{-3}} \right)^{-1} \quad (6)$$

Halo PNs are expected to have a speed of $\gtrsim 100\text{ km s}^{-1}$ through the ISM, while those in the plane move much slower. In addition, close to the galactic plane the density is higher than in the halo. We therefore expect that halo PNs will have $\eta > 1$ and therefore the flow will be adiabatic, while for disk PNs $\eta < 1$ and the flow will be isothermal. Several high quality observations, most notably given in the Atlas published by Tweedy & Kwitter (1996) provide us with a sample of 34 deep images of interacting PNs. In table 1 of Dgani & Soker (1998), we give a list of the objects and their references. Some nebulae have several parallel stripes. We define a nebula as striped only if it has at least two parallel long filaments. All of the 12 striped nebulae are close to the galactic plane ($|z| < 250\text{ pc}$). The fact that the striped nebulae are confined to the galactic plane is compatible with

the prediction that RT will be effective for galactic plane PNs moving in the warm medium, because of magnetic effects, but only for modes perpendicular to the direction of the magnetic field. The result is elongated structures or "RT rolls". For galactic halo interacting PNs the cooling is inefficient and magnetic pressure is negligible. RT will be effective but it will form fingers or blobs and not stripes.

4.4. *Measuring the ISM magnetic field with interacting PNs*

The group of 12 nebulae that show stripes is not homogeneous; some have thick loose stripes, some have thin densely packed stripes. In Dgani (1998) the properties of the striped nebulae and their application to the study of the ISM are explored. In particular, a simple model for the shocked ISM region is used to derive a relation between the distance between adjacent stripes and the strength of the magnetic field of the ISM. If Δz is the average spacing of the stripes in units of the radius of the nebula then (see Dgani 1997, eq. 4):

$$\Delta z \sim v_{A0} \cos \alpha_0 / v_*, \quad (7)$$

where v_{A0} is the Alfven speed in the ISM, α_0 is the angle between the pre-shock magnetic field and the shock front, and v_* is the velocity of the central star. According to the formula nebulae with densely packed stripes either move faster or move in a weaker magnetic field than the ones with thicker looser stripes. Applying this simple formula to several striped nebulae, Dgani shows that information about the strength of the ISM magnetic field in the local neighborhood of these nebulae can be extracted by accurate observations of the members of this group.

Acknowledgement: *It is a pleasure to thank Noam Soker for introducing me to the fascinating subject of interacting nebulae and for a very pleasant collaboration.*

References

- Abell, G. O. 1966, ApJ, 144, 259.
- Borkowski, K. J., & Tsvetanov, Z., & Harrington, J. P. 1993, ApJ Lett., 402, L57.
- Borkowski, K. J., & Sarazin, C. L., & Soker, N. 1990, ApJ, 360, 173 (BSS).
- Brighenti, F. & D'Ercole, A. 1995, MNRAS, 277, 53.
- Dgani, R. 1995, Ann. of the Israel Physical Society, Vol. 11: *Asymmetrical Planetary Nebulae*, eds. A. Harpaz and N. Soker (Haifa, Israel), p. 219.
- Dgani, R. 1998, RevMexAC, 7, 149.
- Dgani, R., & Soker, N. 1994, ApJ, 434, 262.
- Dgani, R., & Soker, N. 1997, ApJ, 495, 337.
- Dopita, M. et. al., 2000, these proceedings.
- Gurzadyan, G. A. 1969, *Planetary Nebulae* (New York : Gordon & Breach), p. 235.

- Hollis, J. M., Van Buren, D., Vogel, S. N., Feibelman, W. A., Jacoby, G. H., & Pedelty, J. A. 1996, *ApJ*, 456, 644.
- Isaacman, R. 1979, *A&A*, 77, 327.
- Jacoby, G. H., & Van de Steene, G. 1995, *AJ*, 110, 1285.
- Jones, T. W., Kang, H., & Tregillis, I. L. 1994, *ApJ*, 432, 194.
- Kaler J. B. 1983, *ApJ*, 271, 188.
- Kerber et. al. 2000, these proceedings.
- McKee, C. F. 1987 in *Spectroscopy of Astrophysical Plasmas*, ed. A. Dalgarno
- Pottash, S. R. 1984, *Planetary Nebulae* (Dordrecht : Reidel).
- Priest, E. R. 1984, *Solar Magnetohydrodynamics*, Reidel Pub. Com., (Dordrecht, holland).
- Raymond, J. C. 1979, *ApJS*, 39, 1.
- Smith, H. 1976, *MNRAS*, 175, 419.
- Soker, N., 1994, *MNRAS*, 270, 774.
- Soker, N., 1996, *ApJ*, 496, 734. .
- Soker, N. & Dgani, R., 1997, *ApJ*, 484, 277.
- Soker, N., Borkowski, K. J., & Sarazin, C. L. 1991, *AJ*, 102, 1381 (SBS).
- Soker, N., Rappaport, S., & Harpaz, A., 1998, *ApJ*. 496, 842.
- Stevens, I. R., Blondin, J. M., & Pollock, A. M, 1992, *ApJ*, 386, 265.
- Tweedy, R. W. 1995, *Ann. of the Israel Physical Society*, Vol. 11: *Asymmetrical Planetary Nebulae*, eds. A. Harpaz and N. Soker (Haifa, Israel), p. 209.
- Tweedy, R. W., & Kwitter, K. B. 1994a, *AJ* 108, 188.
- Tweedy, R. W., & Kwitter, K. B. 1994b, *ApJ* 433, L93.
- Tweedy, R. W., & Kwitter, K. B. 1996, *ApJS* 107, 255.
- Tweedy, R. W., Martos, M. A., & Noriega-Crespo, A. 1995, *ApJ*, 447, 257.
- Tweedy, R. W., & Napiwotzki, R. 1994, *AJ* 108, 978.
- Villaver, E. et. al. these proceedings.
- Weinberger, R. 1989, *A&A Supp.*, 78, 301.
- Zucker, D. B. & Soker, N. 1993, *ApJ*, 408, 579.
- Zucker, D. B. & Soker, N. 1997, *MNRAS*, 289,665.